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A search for X-ray flares in the nuclei of galaxies.

by

John Cunniffe

A Thesis submitted to
The University of Dublin
for the degree of

Doctor in Philosophy

Department of Physics
University of Dublin
Trinity College

November, 2002
Declaration

This thesis has not been submitted as an exercise for a degree at any other University. Except the following listed articles and where otherwise stated, the work described herein has been carried out by the author alone. This thesis may be borrowed or copied upon request with the permission of the Librarian, University of Dublin, Trinity College. The copyright belongs jointly to the University of Dublin, Dunsink Observatory and John Cunniffe.

Long Term X-ray Variability of Galactic Nuclei
J. Cunniffe and E. J. A. Meurs, 1999, I.A.U. Symposium 194, p420,
Frequency of X-ray transients in galactic nuclei

Signature of Author
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Destiny seems to have meant that significant parts of my academic life in the last decade have been strangely affected by one person. Without your illuminating thoughts on politics, onions, Dr. Zeldovich, blind side kidney
punches, cows, real GAA club training, tenants' rights, Latex, and of course how to cook properly, my life in Dunsink would have been a far less colourful one! Many thanks Michael for all your friendship and irrepressable good humour.

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Finally I would like to thank my family for supporting me in so many ways and especially to Vincent and Ronan for all the technical support and for frequently listening to me thinking out loud!
Observations of the giant elliptical galaxy NCG 4552 at ultraviolet wavelengths during the early 1990’s showed sharply peaked emission within ~2pc of its centre which initially brightened and then faded. The features of this flare in emission around the nucleus made it one of the few candidates for a stellar tidal disruption event not identified at X-ray energies. In order to compare its X-ray properties with the existing flare candidates a detailed analysis of the X-ray history of NGC 4552 was carried out. No correlation was found between the UV and X-ray variability and the limits placed on the X-ray output from any faint flare component imply that the mass accreted must have been less than $10^{-5}/\epsilon_{0.1} M_\odot$.

Possible mechanisms for providing such small masses through stellar disruption around the $\sim 4 \times 10^8 M_\odot$ black hole at the centre of the galaxy include tidal stripping of the atmosphere of a giant star and a new scenario in which stellar material from main sequence stars may be ejected after tidal heating. Mechanisms involving variability of an existing low level active nucleus also cannot be ruled out.
During the analysis of the ROSAT High Resolution Imager data for NGC 4552, techniques were developed for assessing the spacecraft attitude error component for faint sources. These methods allowed a determination of the level of broadening in the data but were not sufficiently accurate to de-blur the images.

The ROSAT All-Sky Survey data archive was examined to search for faint flares which may have been missed by previous examinations. The archive data were reprocessed and upper limits fluxes for $\sim 10^5$ LEDA galaxies were determined and compared with detections from the ROSAT pointed observations. None of the $\sim 10^3$ galaxies with pointed detection showed unexpected variability and this determination lowers the upper limit on stellar disruption flaring rate by an order of magnitude from a previous study.
To my parents with love,

for teaching me to ask why.
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Chapter 1

Introduction

In the last couple of decades it has become increasingly clear that most galaxies host massive black holes at their centre. This deduction has been drawn from observations of active galactic nuclei and has been complemented by spectroscopic determinations of stellar and gas dynamics within galaxies.

The strongest argument for a massive central object is provided by infrared observations of stellar proper motions at the centre of the Milky Way (Genzel et al. 1997, Schödel et al. 2002). These orbits indicate a central mass of $(2.6\pm0.2)\times10^6\, M_\odot$ (solar masses) and a mass density of at least $10^{12}\, M_\odot\cdot pc^{-3}$. A cluster of separate stars with such a high density is dynamically unstable due to the rate of collisions between individual stars and would quickly lead to collapsed objects forming at the centre. The timescale for the resulting collapse of the cluster to a central mass is of the order of tens to hundreds of thousands of years. Therefore these cores would have formed very early in
the galaxy's history if indeed they did not pre-date its formation.

For galaxy cores outside our own, stellar orbits are not individually visible but the aggregate stellar velocity dispersion inferred from spectral linewidths allows core masses to be calculated. As high spatial resolution spectroscopy has become available, the radii within which this mass must be confined have been pushed closer to the galaxy centres. These results indicate that many galaxies contain between $10^6$-$10^9 \, M_\odot$ within, at most, a few hundred light years of the centre (Kormendy & Richstone 1995). Although the density lower limits are not as high as for the Milky Way, the collapse timescales are still much shorter than the galaxy lifetimes and the inevitable conclusion is that they contain massive black holes at their centres.

Active galactic nuclei (AGN) were first identified at optical wavelengths by Seyfert (1943) with spectra showing strong broad emission lines. The width of the emission lines imply velocity dispersions of several thousand kilometres per second and this is inferred to be rapid orbital motion of a disk of material around the black hole. As this material orbits, internal friction causes heating of the disk to typically $10^5 K$ and produces the broad emission lines and hot X-ray spectrum. Radiation emitted by these disks can give even stronger evidence of a central black hole by showing the characteristic changes in lineshape due to the strong relativistic effects that occur near the event horizon (Iwasawa et al. 1996). It may also be possible that infalling material crosses the event horizon without efficient radiation processes (Narayan, Yi & Mahadevan 1995) thus making these galactic nuclei appear as very low luminosity objects. Externally, galaxies possessing AGN
are similar in terms of larger structures (disk, bulge, bar components) and stellar luminosity to other galaxies which lack the bright point-like core.

Since they were first launched above the atmosphere at the end of the 1960's, X-ray observatories have detected increasing numbers of optically identified AGN as strong X-ray emitters. This emission is now believed to come from hot plasma at the inner regions of the accretion disk formed as material falls into the central black hole. The various details of the observed features change with the angle we view the central regions but the structures at the centres of most AGN are believed to be largely the same. What is perhaps puzzling is that most galaxies, despite having similar core masses, do not have active nuclei. They do not emit strongly at X-ray wavelengths nor show optical emission lines characteristic of AGN behaviour and it is believed that in these cases no luminous accretion disk has formed around the central black hole. Whether this is due to the absence of fuel falling into the hole, inefficient accretion or the strong absorption of emission around the nucleus remains to be resolved.

Since the ROSAT All-Sky Survey in the soft X-ray region (0.1-2.4keV, 124-5Å) in 1990, several galaxies have been observed which went from emitting at typical AGN luminosities in 1990 to drop in output by a factor of more than one hundred over the following years (Komossa 2001). In their faint states, these galaxies would have been seen as bright normal galaxies (or possibly faint AGNs) but in their bright states they are at the high luminosity end of the AGN distribution. The mechanism proposed for these large changes in output is for a star to pass close enough to the large unfuelled black hole in a
normal galaxy to be tidally disrupted and the resulting debris to fall towards the black hole on a timescale of months to a few years (Rees 1988, Rees 1990). The resulting accretion disk would be expected to emit strongly at X-ray energies for as long as the debris fall back continued after which it would fade quickly to the previous quiescent level.

One interesting example of this transient flaring in a galactic nucleus was observed in 1995 at UV wavelengths in the galaxy NGC 4552 (Renzini et al. 1995). I have carried out an analysis of the variations in X-ray luminosity of this galaxy to search for any correlations with the UV variation. In order to examine the frequency of such soft flares among non-active galaxies I have analysed the ROSAT All-Sky Survey Data and prepared a catalogue of X-ray properties of a large sample of galaxies.

The remainder of this thesis is organised as follows:

Chapter 2 deals with the optics and detectors used in X-ray astronomy and the data processing procedures used with them.

In Chapter 3 I review the current observations and theoretical models of tidal disruption flares.

In Chapter 4 I discuss the detailed X-ray analysis of the UV flare candidate NGC 4552. Particular attention is given to determining the level of soft X-ray variability of the galaxy since its first detection in that band.

In Chapter 5 the results on flaring rates from the ROSAT All-Sky Survey galaxy survey are presented.
CHAPTER 1. INTRODUCTION

In Chapter 6 the implications of this work and future directions are reviewed..

The detailed background to the ROSAT HRI attitude recovery procedures and the also the analysis pipelines for dealing with ROSAT All-Sky Survey data are presented in the Appendices.

Note on Constants and Units: Throughout this thesis I follow the convention in the astronomical literature of using cgs units and also express temperatures and photon energies in terms of keV.

For ease of conversion I provide the SI equivalents here:

<table>
<thead>
<tr>
<th>Quantity</th>
<th>SI</th>
</tr>
</thead>
<tbody>
<tr>
<td>Energy 1erg</td>
<td>$10^{-7}$J</td>
</tr>
<tr>
<td>Luminosity 1erg.s$^{-1}$</td>
<td>$10^{-7}$W</td>
</tr>
<tr>
<td>Flux 1erg.s$^{-1}$.cm$^{-2}$</td>
<td>$10^{-3}$W.m$^{-2}$</td>
</tr>
<tr>
<td>Photon Energy 1keV</td>
<td>$\simeq 1.602 \times 10^{-16}$J</td>
</tr>
<tr>
<td>Temperature 1keV</td>
<td>$\simeq k \times (1.16 \times 10^{7} \text{K})$</td>
</tr>
<tr>
<td>Wavelength 1keV</td>
<td>$\simeq h.c/(12.4 \text{Å})$</td>
</tr>
</tbody>
</table>

I have used $H_0=75 \text{km.s}^{-1}.\text{Mpc}^{-1}$ and $q_0=0$ throughout and where luminosities are quoted from the literature, they have been converted to this model unless otherwise noted.
Chapter 2

Theory

2.1 Origins

The original motivation for considering the disruption of stars near supermassive black holes (SMBHs) was as a possible fuelling mechanism to sustain AGN and QSO light output. It was rapidly realised however that the stellar relaxation processes needed to keep providing stars travelling on near-radial orbits towards the centre of a galaxy were not efficient enough to feed the mass infall rate required by the observed central emission (Hills 1975, Hills 1978, Frank 1978).

Further work was done on the possibility of disruption by a SMBH at the centre of the Milky Way (Lacy, Townes & Hollenbach 1982) but the suggestion (Rees 1988, Rees 1990) that ‘dormant’ SMBHs in normal galaxies could be detected if a star was disrupted and temporary luminous accretion
occurred prompted more theoretical work. This temporary accretion should appear as a short 'flare' lasting a few months to a year, and would be seen principally in the UV and soft X-ray from the inner regions of an accretion disk at temperatures of a few times $10^5$K.

### 2.2 Tidal disruption around a black hole

#### 2.2.1 Physical processes during disruption

Early consideration of the processes going on within the star during disruption focussed on the dissipation of tidal energy in the star during pericentre passage. It was proposed (Lidskii & Ozernoi 1979) that the distortion shock wave passing through the star would lead to the ejection of a heated stellar envelope. This was expected to radiate at $L_{\text{peak}} \sim 10^{41}\text{erg.s}^{-1}$ over timescales of several million seconds (i.e. weeks). Another effect which may occur during close passages is an increase in the rate of nuclear reactions due to the tidal heating of the star (including shock driven processes) (Carter & Luminet 1982, Carter & Luminet 1983, Luminet & Marck 1985, Luminet & Carter 1986). Again this may release enough energy to unbind the star. Neither of these processes is likely to be directly observable but the likelihood of stellar debris fuelling an accretion disk around the BH has led to extensive examination of the processes involved during disruption, debris fallback and accretion. These have been modelled both analytically (Kochanek 1994, Khokhlov & Melia 1996, Loeb & Ulmer 1997, Ulmer, Paczynski &
CHAPTER 2. THEORY


The initial issue to consider is what fraction of the stellar material becomes gravitationally bound during the disruption and what fraction escapes. When a star passes close to a central black hole it will be disrupted if its self gravity is exceeded by the tidal force due to the steep gravity gradient from the black hole. For a Schwarzschild BH this tidal radius is:

\[ r_T = R_\star \left( \frac{\eta^2 M_{BH}}{M_\star} \right)^{\frac{1}{3}}. \]  

(2.1)

where \( \eta \simeq 2.21 \) for a homogenous incompressible body (\( n=0 \) polytrope\(^1\)) and \( \eta \simeq 0.844 \) for an \( n=3 \) polytrope (Sridhar & Tremaine 1992, Diener et al. 1995). In practice, most authors assume \( \eta^{2/3} \simeq 1 \) for calculation of \( r_T \).

For solar type stars \( r_T \) will be close to the Schwarzschild radius, \( r_S = 2GM_{BH}/c^2 \), for \( M_{BH} > 10^8 \) M\(_\odot\) and strong relativistic effects will affect the disruption. For black hole masses much bigger than \( 10^8 \) M\(_\odot\), the disruption will take place within \( r_S \) and no flare would be seen (see Fig 2.1 for the dependance of \( r_T \) on \( M_{BH} \)). For these larger black holes only the outer atmospheres of lower gravitationally bound stars (giants) could be stripped outside \( r_S \) and contribute to luminous accretion (Syer & Ulmer 1999, Di

\(^1\)The polytropic index \( n \) relates density (\( \rho \)) and radius (\( r \)) in stellar models described by the Lane-Emden equation, \( \frac{1}{r^2} d \left( r^2 \frac{dQ}{dr} \right) / dr = -Q^n \), where \( Q = \left( \frac{\rho}{\rho_{central}} \right)^{1/n} \).
Stars passing further from a BH may lose sufficient orbital energy through tidal dissipation to become gravitationally bound and, over repeated pericentric passages, the initially highly eccentric (or open) orbit circularises (Fabian, Pringle & Rees 1975, Press & Teukolsky 1977, Novikov, Pethick & Polnarev 1992). This will ultimately lead to the disruption of the star. Fig 2.6 shows a summary of the zones for advection, disruption and capture.

A solar mass star disrupted passing just inside the disruption radius, $r_T$, experiences a tidal distortion during approach to the pericentre which will lead to a torque on the star. This would ‘spin-up’ the star to close to the
Figure 2.2: The debris trajectories after disruption during a star’s passage within $r_T$ of a SMBH. The range of orbital energies shown is for a solar mass star being disrupted by a $10^6 \, M_\odot$ black hole. The distribution in the binding energies determines the mass fall back rate for the flare. (From Rees 1988)

corotation angular velocity by the the time of closest passage and lead to the higher orbital angular momentum material furthest from the BH becoming unbound and that on the 'inside track' being bound (see Fig 2.2). In a linear perturbation theory treatment of the spin-up process (Press & Teukolsky 1977, Alexander & Kumar 2001, Alexander & Livio 2001), the rotational angular velocity of the star can be described as,

\[
\omega_s \sim \frac{T_2(\eta) v_p}{2\alpha \eta^2 r_p}, \quad \eta = \left(\frac{r_p}{r_T}\right)^{3/2}
\]
where \( v_p \) and \( r_P \) are the velocity and radius at pericentre, \( \alpha \) is the stellar moment of inertia in units of \( M_\star R_\star^2 \) and \( T_2 \) is the second tidal coupling coefficient. A MS star of mass 0.76\( M_\odot \) and radius 0.75\( R_\odot \), passing with pericentric distance equal to the disruption radius, will have \( \alpha \approx 0.07 \) and \( T_2(1) \approx 0.36 \) (Alexander & Kumar 2001). This implies a spin up to \( \omega_s \approx 0.86 \) times the corotation velocity (\( \frac{v_c}{r_p} \)).

Despite the tidally induced transfer of orbital angular momentum to stellar spin, the mean specific binding energy of the stellar material would remain positive (i.e. unbound) but the energies would be spread between,

\[
E_{\text{min}} = -\frac{GM_\star}{r_*} \left[ \left( \frac{M_{BH}}{M_\star} \right)^{1/3} + 1 \right] 
\]

(2.3)

and,

\[
E_{\text{max}} = \frac{GM_\star}{r_*} \left[ \left( \frac{M_{BH}}{M_\star} \right)^{1/3} - 1 \right]
\]

(2.4)

For the assumption of uniformly distributed mass-fractions between these limits, the mass fallback rate of the bound material (i.e. with negative specific binding energy) will be,

\[
\dot{M} \propto M_{BH}^{-1/2} (t/t_i)^{-5/3},
\]

(2.5)

where \( t_i \) is the orbital period of the most tightly bound debris and \( t \) is the time since disruption. Initial simulations (Evans & Kochanek 1989) of the disruption process using smoothed particle hydrodynamics (SPH) showed good agreement with this predicted mass fall back law (Fig 2.3).
Figure 2.3: The modelled mass fall back rate showing predicted long-term $t^{-5/3}$ decay and the highly super-Eddington luminosity expected if the flare was radiating with $\varepsilon=0.1$ around a $10^6 \text{ M}_\odot$ black hole. The orbital period of the debris gives the time from disruption to subsequent fallback. (From Evans & Kochanek 1989)

More recent simulations (also using SPH codes) by Ayal, Livio and Tiran (2000) (Figs 2.4 and 2.5) modelling the dynamics of the debris after disruption show that material returning to pericentre may be heated from shocking and compression and become unbound. This reduces the mass fraction that actually is available for accretion to less than a third of the original mass on bound orbits after disruption (Fig 2.5).

There are some important differences in the case of Kerr BHs to the models...
Figure 2.4: Numerical models of the tidal effects on a star during disruption (top). The lines point towards the SMBH in each figure as the simulation follows the star on its orbit. The lower six images show the particle distribution (left) and density (right) for three simulations of the stream of material returning to pericentre after one orbit. The BH is indicated (+) at the bottom right of the figures. (From Ayal, Livio & Piran (2000) Figs 2 and 5.)
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Figure 2.5: Fraction of material becoming unbound (dot dashed line) or accreting (dashed line) onto the central SMBH during the evolution of the orbiting material. The total mass returning to pericentre (thick line) exceeds the accreted mass (dashed line) due to material becoming unbound in winds created by heating and shocking as the debris streams collide. (From simulations by Ayal, Livio & Piran (2000) Fig 9)

considered above. In particular the differences in orbital angular momentum of the debris are expected to vary significantly depending on the initial orbital angular momentum of the star, $L_*$, the angular momentum of the BH, $J_H$, and the angle between $L_*$ and $J_H$ (Diener et al. 1997). This can have significant effects on the distribution of energies of the debris and consequently on the mass fall back rates. The boundary within which the star is advected before disruption also changes with disruption of solar mass stars possible for certain $L_*$ and $J_H$ around black holes larger than $10^8 M_\odot$ (Beloborodov et al. 1992).
A further observational consequence of a stellar disruption may be seen in the surrounding ISM. If the unbound ejecta collide with the ISM then the conversion of some of the $\sim 10^{52}$ erg (Rees 1988) of kinetic energy may produce (possibly asymmetric) SNR-like structures and this has been invoked as a possible cause of Sgr A East (Khokhlov & Melia 1996).

Figure 2.6: The radial cut-offs for tidal processes for a solar-density star around a massive black hole. Within $r_T$ disruption is possible which may or may not be within the Schwarzschild radius, $r_S$ (dotted curve). Stars passing within $3r_T$ can lose orbital angular momentum and become bound before ultimately being disrupted. For lower density stars, atmospheric stripping may be possible significantly outside $3r_T$ allowing non-advective fueling even for massive BHs with $r_S > r_T$.

2.2.2 Appearance of the flare

Before proceeding to discuss the details of emission from stellar disruption flares, the main features of general accretion structures need to be outlined.
Comprehensive reviews of accretion theory are given in Pringle (1981) and Frank, King & Raine (2002).

To conserve angular momentum material will enter an orbit as it falls towards the centre of a gravitational potential. These orbits will generally cross and dynamical friction will cause heating which will radiate away energy. The interaction will also allow angular momentum to be transported from inner (high J) orbits to outer ones. As the material loses energy the orbit will evolve towards the lowest energy orbit available for its angular momentum at radius, \( r = J^2/GM \) where J is specific angular momentum of the material and M is the central mass. This circularised material will orbit with continued heating and outward angular momentum transport via viscosity in the disk. Dynamical friction is not sufficient to allow the level of accretion observed and the true source of this viscosity in disk responsible for angular momentum transport is still the subject of much debate (Lasota 2001). One candidate for the viscosity is magnetorotational instability in the disk (Balbus & Hawley 1991).

Heating of the disk via shocks at the shear interfaces will cause the temperature to rise. If the disk is geometrically and optically thin (and can therefore cool efficiently) then it will typically radiate with a thermal plasma spectrum characteristic of the abundances, temperatures and densities over the disk. If cooling becomes inefficient (i.e. the disk is no longer optically thin) then the radiation pressure perpendicular to the plane of the disk will cause it to inflate.

There are several reasons why efficient cooling becomes impossible principally
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concerned with the mass infall rate, $\dot{M}$. If $\dot{M}$ is too low then the infalling gas has very low density and therefore a long cooling time, and if $\dot{M}$ is too high then the disk becomes very optically thick and the escape timescales become long compared with the time taken to fall into the black hole. In the latter case the radiation pressure causes an envelope to be supported above the disk (or around the structure entirely) and all radiation emitted from the disk is heavily reprocessed and will radiate as a black body if the outer edge of the 'photosphere' is in thermal equilibrium.

For stellar disruption a significant amount of energy is expected to be released before accretion proper begins. As the streams of material (see Fig 2.2) return to the pericentre and collide with each other, frictional heating of the material will tend to form a hotspot which will radiate away significant orbital energy as the streams circularise. The rate of energy release will be dominated by the mass fallback rate and this part of the emission may emit above the Eddington limit for a considerable time as the radiation pressure over the whole disk and surrounding is insufficient to stop the infalling high density debris streams. During this 'super-Eddington' phase some of the returning material will become unbound and driven away in a heated wind (Ulmer 1999).

Apart from the emission due to the shock heating and large cooling area of the stellar material during the fall-back phase, the dominant observable feature of a disruption event is expected to be the long term accretion of the bound debris. Initial models (Rees 1988, Rees 1990, and references therein) assumed that half of the mass of the disrupted star would be bound to the BH
and fall back as $t^{-5/3}$ (Eqn 2.5). As the most tightly bound material returns to pericentre it will be highly shocked and would rapidly circularise (at twice the pericentric radius to conserve angular momentum) in a viscous accretion torus. The total energy release during mass fall back is $G M_{BH} \Delta M / (2 r_c)$, where $r_c$ is the circularisation radius.

The torus material would then spread to both larger and smaller radii and begin to accrete onto the BH at $\dot{M} \propto t^{-1.2}$ (Cannizzo, Lee & Goodman 1990). The luminosity during this viscous accretion stage is expected to be much lower than the mass fall back stage and the accretion timescale much longer ($\sim 100$ yrs) (Li, Narayan & Menou 2002).

Depending on the viewing angle of the material accreting onto the BH, hotter or cooler parts of the accretion disk would be seen. Models for edge-on and face-on views for a $10^7 M_\odot$ are shown in Fig 2.7.

An alternative model involves the emission from the accretion disk creating a radiation supported envelope around the black hole radiating at close to its Eddington luminosity. Since $L_{edd}$ is a function of black hole mass,

$$L_{edd} = \frac{4 \pi G m_p M_{BH} C}{\sigma_T} \simeq 1.3 \times 10^{38} \frac{M_{BH}}{M_\odot} \ \text{erg.s}^{-1}$$ (2.6)

where $m_p$ is the mass of the proton and $\sigma_T$ is the Thompson scattering cross section. Since this envelope would be radiating at close to $L_{Edd}$, it has been proposed as a measurement for black hole masses (Ulmer 1999). The key observable feature of the envelope would be that reprocessing of the radiation would mean that its surface would radiate at a temperature of $T \approx 10^4$ K rather
Figure 2.7: Spectra for models of a flare around a $10^7 M_\odot$ black hole (flux for $D=100\text{Mpc}$). The solid lines show calculated spectra from an optically and geometrically thick disk as viewed from face-on and edge-on. The former is much brighter and hotter because a large fraction of the energy is radiated in the funnel. A black body curve (dashed line) with $T = 10^5$ (kT$\approx0.01\text{keV}$) is shown for comparison (From Ulmer 1999).

than the $10^5 - 6K$ expected if the accretion disk was directly visible (Loeb & Ulmer 1997, Ulmer, Paczynski & Goodman 1998, Ulmer 1999).

A separate observable effect of the tidal stripping of the atmosphere of a giant star is the possibility of leaving a hot He star as an observable remnant. This has been proposed as a candidate for the $L_X \approx 10^{37}\text{erg.s}^{-1}$ super-soft source at the centre of M31 (Di Stefano et al. 2001).
2.3 Expected disruption rates

For a star to be disrupted its pericentre distance must be less than $r_T$. The standard approach to determining the disruption rate is to consider which orbits will have pericentres within $r_T$ (Frank & Rees 1976, Lightman & Shapiro 1977, Rees 1988, Magorrian & Tremaine 1999, Syer & Ulmer 1999, Freitag & Benz 2002). This 'loss cone' of orbits (see Fig 2.8) is described by,

$$\theta_{lc} \simeq 2 \frac{r_T GM_{BH}}{r^2 \sigma^2}$$  (2.7)

where $\sigma$ is the isotropic stellar velocity dispersion at radius $r$. The loss cone is usually very small with typical values of $\theta_{lc}^2$ of $10^{-5}$.

Figure 2.8: Loss cone orbits around a black hole are those which have pericentres within $r_T$ (or equivalently, angular momentum $J < J_{lc}$) (from Freitag & Benz 2002).

For solar mass stars with an isotropic velocity distribution, the expected rate
at which stars would pass within a distance $r_{\text{min}}$ of the centre is given by,

$$\sim 10^{-4} \left( \frac{M_{\text{BH}}}{10^6 M_\odot} \right)^{4/3} \left( \frac{N_*}{10^5 \text{pc}^{-3}} \right) \left( \frac{\sigma}{100 \text{km.s}^{-1}} \right)^{-1} \left( \frac{r_{\text{min}}}{r_T} \right) \text{yr}^{-1}$$  \hspace{1cm} (2.8)

where $N_*$ is the stellar density, and $\sigma$ the velocity dispersion of stars in the centre of the galaxy (Lightman & Shapiro 1977, Rees 1988). This rate does not reflect the contribution from stars passing outside $r_T$ which may be subject to tidal capture at radii up to $3r_T$ (Novikov, Pethick & Polnarev 1992) which would increase the $\left( \frac{r_{\text{min}}}{r_T} \right)$ term by 3.

Stars on loss cone orbits will be disrupted within one orbital period and therefore the actual rate will be that at which the loss cone orbits are re-filled. The dominant process in repopulating the loss cone is two body stellar scattering. The simplifying assumption is Eqn 2.8 have been broadened in more recent work (Magorrian & Tremaine 1999, Syer & Ulmer 1999, Freitag & Benz 2002) with calculations of scattering rates for various estimates for stellar density and velocity distribution (both isotropic and flattened). These lead to estimates in the range $10^{-4}$-$10^{-5}\text{yr}^{-1}$ for solar mass stars and $10^{-4}$-$10^{-9}\text{yr}^{-1}$ for disruption of giants around black holes bigger than $10^8 M_\odot$. These values depend very strongly on the assumed parameters.

For the Keplerian potentials close to the central BH it has been shown (Rauch & Tremaine 1996, Rauch & Ingalls 1998) that resonant relaxation will allow an increase in angular momentum transport by a factor of a few. However, this effect only operates close to the BH and is unlikely to significantly increase the overall population of stars being consumed.
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Another mechanism which may affect the rate at which stars can enter loss cone orbits is through disturbed velocity structures within a galaxy. At least one of the observed flare candidates (NGC 5905, see next section) has multiple bar structures (Wozniak et al. 1995, Friedli et al. 1996). The possibility of stable triaxial velocity distributions (Poon & Merritt 2002) in the nuclei of galaxies is being explored and this would lead to different rates for loss cone filling processes in comparison with standard distribution models (Freitag & Benz 2002).

There is also a contribution to loss cone repopulation from stellar evolution due to the dependance of \( r_T \) on stellar density. In particular it is important for old MS stars just outside the loss cone (Syer & Ulmer 1999). As they become giants they can find themselves within \( \theta_c \) which increases approximately as \( R^{1/2} \) (Eqn 2.1 and Eqn 2.7).

2.4 Observed X-ray flares

During analysis of the ROSAT All-Sky Survey (RASS) data, several bright soft objects were found which were seen to be much fainter in follow up observations. These discoveries led to an extensive search of the X-ray archive for objects showing large changes in brightness or in spectral shape (indicating significant change in one emission/absorption component.)
2.4.1 Flares in active galaxies

The first extragalactic object which was observed with a large amplitude change in soft X-ray brightness was the Seyfert 1 galaxy 1E1615+061 (see Table 2.1 for summary of flaring objects). Having been observed by the SAS 3 and HEAO 1 satellites as a bright soft source (Pravdo et al. 1981), it was observed to drop in brightness by a factor of ~100 between the observation by Einstein in February 1980 and that by EXOSAT in July 1985. The spectrum also flattened considerably between the two observations with a powerlaw fit changing from $\Gamma \sim -4$ to $\Gamma \sim -2$ ($f_\nu \propto \nu^\Gamma$) (Piro et al. 1988, Piro et al. 1997, Guainazzi et al. 1998b).

Two of the bright soft sources detected during the RASS were IC 3599 and WPVS007. IC 3599 (Zw59-34, RXJ 1237+264) was observed in a ROSAT PSPC pointing one year later a factor of 80 lower in count rate (Grupe et al. 1995b, Brandt, Pounds & Fink 1995) and over the following 18 months it fell by a further factor of ~3. Over this time the spectrum also became somewhat softer with reducing count rate.

Similarly, the narrow-line Seyfert 1 (NLSy1) galaxy, WPVS007 (RXJ 0039.3-5117), was observed to drop by a factor 400 between the RASS and a PSPC pointing 3 years later (Grupe et al. 1995b). There is no strong evidence for spectral change between the two observations with the hardness ratio (Soft=0.1-0.4keV, Hard=0.4-2.4keV, see Eqn 3.9) changing from -0.94±0.06 to -0.86±0.14.

Another object showing significant change in spectral shape is the NLSy1
galaxy RXJ0134-4258 (Grupe et al. 1999, Komossa et al. 2000) which changed from an \( \Gamma = -4.4 \) to \(-2.2\) over two years between the RASS and pointed observations. There was only a moderate (\(~20\%\)) reduction in count rate during the same period. Such a small drop in flux given the large change in spectral slope has lead to suggestions that this may be due to a variable warm absorber (Komossa & Fink 1997) rather than reduction in a soft excess component (Komossa et al. 2000).

Other active galaxies have shown significant variability in their soft excess component including 1H0419-577 (Guainazzi et al. 1998a) which showed a marked soft X-ray excess in a 1992 PSPC observation compared with later ASCA and SAX observations when it had dropped by a factor of \(~6\). Variation in the soft component is also visible in the Seyfert 1 galaxy RX J2248-511 (Breeveld, Puchnarewicz & Otani 2001). As shown in Fig 2.9, the soft X-ray flux seen in the PSPC in 1993 is not consistent with the ASCA spectrum from 1997. The soft X-ray and optical variability are not taken simultaneously but it is proposed that this galaxy may be part of a class of 'Big Blue Bump Variables' in which the ultra-soft X-ray component is the hard tail of the bump.

### 2.4.2 Flares in non-active galaxies

The objects described above differ from the five objects listed in the lower section of Table 2.1 in that the latter have no optical emission features associated with nuclear activity.
Figure 2.9: The Sy 1 galaxy RE J2248-511 from Breeveld et al 2001 (Fig 4) showing variable optical spectrum with significant blue enhancement in 1991 which is not present one year later. The soft X-ray excess (particularly <0.3keV) seen in the 1993 PSPC observation is not seen 4 years later by ASCA where 0.5-10keV is consistent with a flat power-law. and the soft 0.5-2.0keV is significantly lower than the PSPC.
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<table>
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<tr>
<th>Name</th>
<th>$z$</th>
<th>Type</th>
<th>Morph Type</th>
<th>$L_X(max)$</th>
<th>$kT_{bb}$</th>
<th>$L_X(max)$/$L_X(min)$</th>
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<td>1E1615+061</td>
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<td>Sy 1</td>
<td>-</td>
<td>$5 \times 10^{45}$</td>
<td>$\sim 0.02$</td>
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<td>0.028</td>
<td>NLSy 1</td>
<td>-</td>
<td>$1.5 \times 10^{44}$</td>
<td>0.02</td>
<td>400</td>
</tr>
<tr>
<td>IC 3599</td>
<td>0.021</td>
<td>Sy</td>
<td>SOpec</td>
<td>$5 \times 10^{43}$</td>
<td>0.09</td>
<td>200</td>
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<td>NGC 5905</td>
<td>0.011</td>
<td>H II</td>
<td>SB(r)b</td>
<td>$5 \times 10^{42}$</td>
<td>0.05</td>
<td>200</td>
</tr>
<tr>
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<td>-</td>
<td>Pair</td>
<td>$9 \times 10^{43}$</td>
<td>0.06</td>
<td>&gt;20</td>
</tr>
<tr>
<td>RXJ1331-3243</td>
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<td>-</td>
<td>-</td>
<td>$\sim 10^{44}$</td>
<td></td>
<td></td>
</tr>
<tr>
<td>RXJ1624+7554</td>
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<td>-</td>
<td>-</td>
<td>$2 \times 10^{43}$</td>
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<td>200</td>
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<td>-</td>
<td>-</td>
<td>$8 \times 10^{44}$</td>
<td>0.04</td>
<td>150</td>
</tr>
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Table 2.1: Objects showing large amplitude X-ray variability. The upper section shows the three objects with active optical classifications and the lower section shows those without optical indications of activity (HII emission or no optical emission lines).

NGC 5905 (Bade, Komossa & Dahlem 1996, Komossa & Bade 1999) dropped from a count rate of 0.4cts/sec during RASS to less than 0.09cts/sec five months later and less than 0.004cts/sec eighteen months after the RASS detection. The count rate actually rose by a factor $\sim 3$ during four days of the RASS observation. During this time the spectrum was very soft with a blackbody temperature of $kT_{bb}=0.06$keV. This hardened from a RASS hardness ratio of $-0.86 \pm 0.03$ (Soft=0.1-0.5keV, Hard=0.5-2.4keV, see Eqn 3.12) to $-0.65 \pm 0.11$ during a PSPC pointing 3 years later. The lightcurve shown in Fig 2.10 shows the drop in flux observed by ROSAT. However, it should be noted that the lightcurve sampling for NGC 5905 is very sparse and while it is clear that a large drop in output has occured, it is not clear that is follows a monotonic decay law as suggested by the simple theoretical arguments in §2.2.1.

Modelling the flare in NGC 5905 Li, Narayan & Menou (2002) fit the data
with the expected $t^{-3/2}$ decay law. Integrating the emission over the total lifetime of the flare, they find the total emitted energy to be $\sim (4.5 \pm 0.9) \times 10^{49} \text{ erg}$ and the associated accreted mass $2.5 \times 10^{-4} \quad M_\odot \left(\frac{\epsilon}{0.1}\right)^{-1}$, where $\epsilon$ is the standard accretion efficiency parameter ($E = \epsilon M c^2$). They conclude that this mass could have been provided by either partial stripping of the outer layers of a low-mass main sequence star or as the disruption of a brown dwarf or a giant planet (Li, Narayan & Menou 2002).

Similar to NGC 5905, the remaining four galaxies in Table 2.1 show no indications of optical emission lines. The optically inactive galaxy pair RXJ 1242-1119 (Komossa & Greiner 1999), RXJ 1331-3243 (Reiprich & Greiner 2001), RXJ 1420+5334 (Greiner et al. 2000), and RXJ 1624+7554 (Grupe,
Thomas & Leighly 1999) all show strong variability between the ROSAT All-Sky Survey and pointed ROSAT PSPC observations. The lightcurve for RXJ 1420+5334 is shown superimposed on the NCG 5905 curve in Fig 2.10. The decay rates for these objects are not well constrained with in each one detected during outburst and a faint detection or upper limit no earlier than one year later.

2.4.3 The Milky Way

A candidate for transient AGN activity in the recent past is our own galaxy. The nucleus around Sgr A* has $L_X$ (2-10keV) $\sim 10^{35}$erg.s$^{-1}$. However, at a distance of 300lyrs away from Sgr A*, the Sgr B molecular cloud region appears much brighter than can be explained by reflection from the present nucleus (Koyama et al. 1996). The proposal is that Sgr B could be reflecting light emitted when the nucleus had X-ray emission of $\sim 2 \times 10^{35}$erg.s$^{-1}$ in the past. This would most likely have been a short-term state since other clouds at different distances from Sgr A* do not show the same reflected luminosities.

This light echo method of tracing back the accretion history of the nucleus is not as easily applicable to other galaxies due to the faintness of the reflections from the clouds relative to the surrounding material. Also, though such image scales (300lyr $\approx$ 30” at the distance of M31) can be easily imaged by today’s X-ray imagers, determining the reflection geometry of the molecular clouds would prove more difficult.
2.4.4 Optical indication of stellar disruption

A further indication of transient accretion is the variable, broad, double-peaked H$_\alpha$ lines in the nucleus of the Seyfert/LINER galaxy NGC 1097. These features have varied significantly over several years and have been argued as evidence for tidal disruption of a star to provide the transient disk structures observed (Eracleous et al. 1995, Storchi-Bergmann et al. 1995, Storchi-Bergmann et al. 1997, Storchi-Bergmann et al. 2002).

Tidal disruption has also been invoked as a possible explanation to explain the transient increase in strength and linewidth of the He II line from the Seyfert NGC 5548 (Petersen & Ferland 1986).

2.5 Observed UV flares

In 1993 the galaxy NGC 4552 was observed to have a sharp 'spike' in UV images taken by the HST Faint Object Camera (FOC) (Renzini et al. 1995). This feature was found to be $\sim 4.5$ times brighter than the central peak in a similar 1991 observation. Follow-up observations with the post-COSPAR FOC showed a drop by a factor of 2 from the 1993 level (Cappellari et al. 1999). The details of this flare and follow-up at X-ray energies is discussed extensively in Chapter 5.

The search for other UV flares of this type are limited essentially to the FOC archive. The high spatial resolution is required to isolate the faint peak
against the strong emission from the rest of the galaxy. The NGC 4552 flare would have gone undetected in an imager with low spatial resolution due to the brightness of the underlying galaxy. Capellari et al. have searched the FOC archive for similar candidates to NCG 4552. Of the two they discuss, NGC 2681 has a central excess but it remains at a constant brightness and can be explained by a central star cluster and for NGC 1399 the FOC image appears smooth with no clear central spike.

CHAPTER 2. THEORY

Much of this chapter focuses on determining the variability of sources by comparing observations with different instruments. In order to understand the factors affecting such observations and possible systematic differences it may be prudent to pre-process the resulting data. I will discuss the stability of different detectors used in X-ray astronomy. I also consider some of the potential calibration issues. I discuss problems which might be encountered in the analysis of data from detectors.
Chapter 3

X-ray Instrumentation

3.1 Introduction

Much of this thesis focusses on determining the variability of sources by comparing detections with different instruments. In order to understand the factors affecting each observation and possible systematic differences which may be present in the resulting data, I will discuss the standard optics and detectors used in X-ray astronomy. I also consider some of the instrumental calibration uncertainties and problems which affect calculation of spectra and fluxes from data.
3.2 X-ray Optics

In the X-ray energy range of 0.1 keV and 100 keV the complex refractive index,

\[ n_c = 1 - \delta - i\beta \]  \hspace{1cm} (3.1)

for all materials has very small decrements, \( \delta \) and \( \beta \). This implies very weak refraction and would require significant thickness of material in a lens. However, the absorption coefficient is large enough in all cases to completely absorb the incident radiation in any practical lens design and so rules out the use of transmissive optics for X-rays.

Since X-rays are totally absorbed near normal incidence, conventional reflecting optics are not possible. Instead, since the real part of the refractive index \((1 - \delta)\) is negative, total external reflection (TER) will occur at shallow reflection angles. For reflection in a vacuum, the critical angle for TER is,

\[ \theta_{TER} = \frac{\pi}{2} - \delta \text{ radians} \]  \hspace{1cm} (3.2)

For rays reflecting more steeply than \( \theta_{TER} \), the reflection coefficient will drop below 1. Also, since \( \delta \) reduces rapidly with increasing energy, the reflectivity will drop also. The effects of incident angle and energy can be seen in the plots of effective mirror area for the ASCA mirror assembly (see Fig 3.8).

Using glancing reflections to form images is more complicated than normal incidence optics since it takes two reflections to bring rays to a focus over any useful field of view. The first practical imaging X-ray mirror geometry
was proposed by Kirkpatrick & Baez (1948) using parabolas of translation reflecting in orthogonal planes. The first axially symmetric (centred) design was described by Wolter (1952) involving reflection from the inner faces of first a paraboloid of rotation and then a hyperboloid of rotation (see Fig 3.1), which has remained the standard design (see Fig 3.1) in all X-ray observatories to date. This mirror pair will only image light entering an annulus around the axis so in practice several sets of these pairs are nested inside one another with a common focus to increase the collecting area. The Wolter type II and III arrangements involve reflections from the inner face of the paraboloid and then the outer face of a hyperboloid placed closer to the focus (type II) or the inner face of the paraboloid and an the outer face of an ellipsoid (type III). Both of the latter design are more complicated to manufacture and mount and do not allow close nesting of mirror shell so are not used and the Wolter type I design has remained the standard design in all X-ray observatories to date.

Longer focal length designs allow shallower glancing angles, and therefore have higher reflectivities extending to higher energies but also smaller effective areas. To maintain collecting efficiency within the total aperture, more nested mirror sets are needed. However, if these are placed too close together there will be a significantly increase in vignetting (i.e. rays entering part of the annulus do not have a clear path to the focus.)

One important feature of Wolter type I optics is that the focal plane is spherical. This is more pronounced in shorter focal length systems. Where a flat detector placed at the focal 'plane' (or more strictly speaking at the on-
Figure 3.1: The Wolter type I arrangement (top) showing the two optical surfaces (dotted lines) and the two actual mirror sections (solid lines). The rays (dashed lines) strike the paraboloid and then hyperboloid on the mirror face closest to the optical axis. (From Willingale 1999). The Wolter type II arrangement (not shown here) involves a reflection from the inner face of the paraboloid and the outer face of a hyperboloid placed closer to the focus. The ROSAT satellite schematic (bottom) shows the same optical arrangement for X-ray imaging. The rays pass through a collimator to absorb stray light and strike four confocal mirror pairs with an effective focal length of 2.4m, before reaching the focal plane assembly. The three detectors (2 PSPCs and 1 HRI) can be rotated into the X-ray focus. (ROSAT User’s Handbook)
axis focal point), only the central region will be in focus. The outer regions of the detector will be illuminated by a diverging cone which has passed through focus before reaching the detector. This is clearly seen in the Einstein and ROSAT imagers.

The limitations on the useful field of view come from the reduced performance of the mirrors as the off-axis angle increases. Rays from an off-axis angle strike the mirrors at a range of angles with a reduction in effective area. Beyond a certain angle, determined by the length of the mirror sections and the spacing between shells, rays entering part of the annulus may be unable to reach the focal surface leading to further loss of throughput.

Several factors contribute to size of the point spread function (PSF) of X-ray mirrors:

- accuracy of alignment of the optical elements, in particular all mirror pairs have a common focus over the field,
- scattering from the reflecting surface due to micro-roughness in the mirror coating,
- scattering and diffraction from mirror shell support structures,
- deviations of detectors from the spherical focal surface.
CHAPTER 3. X-RAY INSTRUMENTATION

3.3 X-ray Detectors

The original detectors used in X-ray astronomy were Geiger-Müller tubes and ionisation chambers which had very little energy resolution. Since then the sensitivity, spatial and spectral resolution of X-ray detectors have improved significantly. In the present generation of X-ray observatories several instrument designs have been used.

3.3.1 Proportional Counters

Proportional counters involve a gas absorber between multi-wire electrodes with the gas chosen to have a suitable absorption cross-section for the desired energy range. Photoelectric absorption varies as $Z^4/E^{9/3}$ so low Z materials are used for the entrance window (usually carbon or plastic) with a thin metal layer to absorb optical/UV photons and allow control of electrostatic potential. Common absorber gases for soft X-ray detection are Argon (Z=18) and Xenon (Z=54) since they are stable, have a suitable absorption cross-sections and have low electron affinity.

An X-ray photon entering the detector is absorbed by the gas and the photon energy goes to generating a number of electron-ion pairs approximately proportional to the initial photon energy. These pairs are then separated and accelerated in an electric field, creating more electron-ion pairs (mainly due to the electrons) and the resulting amplified current pulse is measured at the collecting electrodes. In order to prevent UV photons emitted after
collisional excitation creating more cascades, a small amount of an organic gas (e.g. CH$_4$ or CO$_2$) may also be added to the absorber gas as a quenching agent (Delaney & Finch 1992). The arrival position on the detector is determined by having the anodes and cathodes in the form of a multi-wire grid. The wire (or wires) on which the charge cloud is collected is then used to calculate the X-Y arrival position. The strength of the collected current pulse is determined by a pulse height analyser (PHA) and used to calculate the energy of the original X-ray photon. This basic arrangement is known as the Position Sensitive Proportional Counter (PSPC).

**Spatial resolution**

The spatial resolution of a PSPC is a function of the spacing of the wire collector grids and depends on the uncertainty of the spatial extent of the charge cloud created in the initial ionisation and subsequent acceleration. Positions between wires can be interpolated by measuring the fraction of charge arriving on neighbouring wires. This process is complicated at the low energy end of the spectrum where fewer electron-ion pairs are created leading to more chance of the fraction on one of the wires falling below the detection threshold.

**Spectral resolution**

In general, the gas may be characterised as requiring an effective energy, W (which may be energy dependent), per charge pair created. So for an X-ray
photon of energy, $E$, the number of electron-ion pairs created will be:

$$N = \frac{E}{W}$$

(3.3)

The variance on $N$ will determine the uncertainty in the measured photon energy. In general this measured uncertainty will be smaller than that expected from random Poissonian statistics by a factor $F$, the Fano factor. So the measured variance is:

$$\sigma^2 = F.N$$

(3.4)

If we assume a Gaussian approximation for the measured pulse height distribution then the FWHM energy resolution is:

$$\Delta E = 2.36 \left[ \frac{F.E}{W} \right]^{1/2}$$

(3.5)

For the avalanche PSPC described above, there is in addition to the uncertainty (Eq. 3.3) in the initial charge cloud creation, an uncertainty in the gain process. These factors lead to an effective uncertainty,

$$\Delta E = A.E^{1/2}$$

(3.6)

where $A$ is an (approximately) energy independent term containing the Fano factor, work function and gain uncertainties. For real devices these factors will have slight energy dependancies but the term in Eq. 3.4 will dominate.

The performance of the detector in terms of the spectroscopic resolution, $R$,
is given by,

$$\frac{\Delta E}{E} = \frac{1}{R} = A.E^{-\frac{1}{2}}$$  \hspace{1cm} (3.7)

This is the figure of merit for any detector whose initial count statistics are proportional to the photon energy. This includes PSPCs, gas scintillation proportional counters, and CCDs. Typical values of $A$ in Eq. 3.7 for PSPCs for is $\sim 0.35$.

### 3.3.2 Gas Scintillation Proportional Counters

A significant improvement on the basic PSPC design is the Gas Scintillation Proportional Counter. In this design the incident X-ray photon is absorbed as before by the detector gas generating electron-ion pairs. These are then accelerated by a relatively low potential which leads the electrons to excite but not ionise the gas atoms. The resulting UV emission is detected by an imaging phototube and the resulting intensity is proportional to the original X-ray energy. Since the gain uncertainty of the PSPC has been eliminated by detecting the UV photons directly, a significant reduction in the value of $A$ in Eq. 3.7 is achieved. A typical value of $A$ for GSPCs is 0.14.

### 3.3.3 CCD detectors

In silicon CCDs, the photoelectric absorption of an X-ray produces approximately one electron-hole pair for every 3.7eV of photon energy. This charge is confined in the volume in which it was created until the whole imaging
array is read out through a series of gates which are clocked in sequence. The normal sequence is to transfer the entire array of pixels to a frame store area and then clock the individual rows, and pixels within each row through an amplifier and by monitoring the current, the charge from each pixel can be determined. The principal uncertainties in determining the energy of the incident photon are noise in the readout electronics and variation in charge transfer efficiency (CTE) across the surface of the chip.

Overall typical CCDs for X-ray astronomy have values of $A$ (Eq. 3.7) of $\sim 0.045$. One factor negatively affecting CCD performance (relative to proportional counters or MCPs) is photon pile-up at high count rates. Here multiple photon events arriving at the same pixel lead to the total charge deposited being read out as the total signal and a summed energy event being registered. This effect can be taken into account in simulations and can be reduced by increasing the CCD readout rate.

The dominant factor affecting CCD spectral resolution in orbit is radiation damage. Cosmic rays passing through the detector lead to damage to the fabric of the detector and readout electronics resulting in individual pixels becoming unresponsive and changes in the CTE over time. These effects increase the effective value of $A$ (Eq. 3.7).

The spatial resolution is primarily determined by the pixel size though charge leakage to neighboring pixels may degrade this somewhat.
3.3.4 Micro Channel Plates

Micro Channel Plates (MCPs) offer high spatial resolution through the use of narrow 'pores' in a glass plate. With diameters of a few microns and channel lengths of the plate thickness (~1mm) these can be packed very densely on the face of the plate. The pore walls are coated with a resistive photocathode material and a potential is applied between the two faces of the plate. An X-ray striking the top of one of the pores will eject an electron which will be accelerated down the length of the channel ejecting more electrons at each collision with the wall. The net gain of this process leads to $10^6$-$10^7$ electrons being produced.

The high spatial density of the pores in the imaging surface means that when an X-ray strikes the detector, the resulting electron charge cloud will be confined in a small volume. This allows a closely spaced multi-wire readout system to discriminate the position of the cloud to within a few microns. In comparison with a PSPC, the major change is the reduction in the size of the amplified charge cloud. However, the large gain and uncertainty in the first few amplification stages as electrons progress through each MCP channel mean that the final integrated charge detected at the grid is not strongly dependent on the X-ray energy. This greatly limits the use of MCPs in imaging spectroscopy and at best an indication of the hardness of the source spectrum is possible.
3.4 X-ray Observatories

<table>
<thead>
<tr>
<th>Observatory</th>
<th>Operation</th>
<th>Detector Type</th>
<th>$\Delta X[\text{&quot;}]$</th>
<th>$\Delta E/E[%]$</th>
</tr>
</thead>
<tbody>
<tr>
<td>Einstein</td>
<td>11/1978 - 4/1981</td>
<td>IPC PSPC</td>
<td>60&quot;</td>
<td>~1</td>
</tr>
<tr>
<td>EXOSAT</td>
<td>5/1983 - 4/1986</td>
<td>LE/CMA MCP</td>
<td>18&quot;</td>
<td>-</td>
</tr>
<tr>
<td>ROSAT</td>
<td>6/1990 - 2/1999</td>
<td>PSPC HRI</td>
<td>15&quot; 5&quot;</td>
<td>0.43</td>
</tr>
<tr>
<td>ASCA</td>
<td>2/1993 - 3/2001</td>
<td>SIS CCD GIS GSPC</td>
<td>3' 3' 0.02† 0.08†</td>
<td>0.09-0.04</td>
</tr>
<tr>
<td>Chandra</td>
<td>7/1999 - present</td>
<td>ACIS-S CCD</td>
<td>~2&quot;</td>
<td></td>
</tr>
</tbody>
</table>

Table 3.1: Summary of the X-ray observatory characteristics. Note (1) the effective spatial resolution of XRT+detector at ~1keV. Note (2) the energy resolution at 1keV (except ASCA at 5.9keV).

3.4.1 Einstein

The Einstein (HEAO-2) satellite (Giacconi, et al. 1979) was launched in November 1978 and carried out the first soft X-ray imaging of the sky. It contained a Monitor Proportional Counter (non-imaging PC), High Resolution Imager (MCP imager), Grating Spectrometer (diffraction gratings with PSPC detector), Solid State Spectrometer (non-imaging Si(Li) crystal), and an Imaging Proportional Counter (IPC) which I will describe in more detail.

The IPC imager was a standard gas filled (84% Ar, 6% Xe, 10% CO$_2$) proportional counter operating with a gas gain of $\sim 10^5$ and read out by crossed anode and cathode grids. Within the central square degree of the detector the readout wire spacing allowed an event to be localised to 1mm which corresponds to 1'. In the outer part of the detector the resolution was somewhat
poorer than this out the full field of view of 75'× 75'. The effective spatial resolution was 1' at 1.5keV degrading at softer energies to 2' at 0.28keV.

![Energy vs Effective Area](image)

Figure 3.2: The effective area for the Einstein IPC-A with the 2.4μm polypropylene+Lexan entrance window. The data points indicate the values at ground calibration (From Giacconi (1979)).

The energy resolution $\Delta E/E$ was $\sim 1$ at 1.5keV and above, and $\sim 2$ at 0.28keV. The effective area for the IPC is shown in Fig 3.2 showing the strong carbon-K edge absorption (0.28keV) from the entrance window. Due to the reflectivity of the Einstein XRT, the effective upper energy cut-off is 3.5keV.
EXOSAT was launched in May 1983 and continued to observe until April 1986. It carried a range of instruments covering the spectrum from 0.05-50keV.

The instruments considered in this thesis from EXOSAT are the low energy (LE) telescopes (0.05-2.5keV) used with the Channel Multiplier Array (CMA) micro channel plate imagers. The LE telescopes are standard Wolter type 1 designs. The CMA detectors are normal MCP detectors but are read out by a four-electrode resistive plate anode which determines the arrival position of the charge cloud from the relative signals on the electrodes. The on-axis half energy width of the PSF is 24", which blurs to 4' at 1° off-axis. Vignetting in the telescopes reduces the effective area to 45% of its peak value at 1° off-axis.

The CMAs have no intrinsic energy resolution and rely on filters with different bandpasses being rotated into the beam to give basic filter photometry of sources (de Korte et al 1981, EXOSAT Observer's Guide Part 2). The spectral response of the different filters used is shown in Fig 3.2. The effective area of the LE/CMA instrument for the 3000Å Lexan (3Lx) filter is ~10cm² on-axis at its peak transmission at 0.15keV. The Aluminium/Parylene (Al/P) response peaks at 0.07keV with an effective area of ~6cm².
CHAPTER 3. X-RAY INSTRUMENTATION

3.4.3 ROSAT - Position Sensitive Proportional Counter

The Röntgensatellit (ROSAT) was launched in June 1990 and continued to operate until February 1999. It contained a 2.4m X-ray telescope (XRT) consisting of four Wolter type 1 mirror pairs with a rotating optical bench at the focus to select the imaging detector (see Fig 3.1). The satellite also contained a separate extreme UV Wide Field Camera. The ROSAT XRT Point Spread Function (PSF) is broadened by a number of effects. The fraction of photons suffering scattering from the microroughness of the mirrors is measured as $0.075E^{1.43}$. If the grazing reflection angle remains constant,
the fraction should increase as $E^2$ but the effective reflection angle averaged over all the mirror sets is not constant hence the lower energy dependance. The off-axis blur circle of the XRT was modelled as a Gaussian with

$$
\sigma(\theta) = \sqrt{(108.7E^{-0.888} + 1.21E^6 + 0.219 \cdot \theta^{2.848}) \text{ arcsec}} \quad (3.8)
$$

for off-axis angle $\theta$ in units of arcminutes (Boese 2000).

The primary X-ray imager, the Position Sensitive Proportional Counter (PSPC) (Pfeffermann et al. 1988) is a standard PSPC design (see Fig 3.4) with a gas composition of 65% argon, 20% xenon and 5% methane which provides a gain factor of $\sim5 \times 10^4$. The grids are arranged as shown in Fig 3.4 with anode spacings of 1.5mm (A1) and 2mm (A2) and cathode spacing of 0.5mm (K1,K2). A set of anti-coincidence grids at the base of the detector will register a pulse from cosmic rays entering the detector, allowing these events to be rejected.

The ROSAT XRT focusses the thin light cones (of half angle $8^\circ$) from each of the mirror sets onto the detector. This leads to some spreading of the light cone leads to an decrease in the spatial resolution due to absorption happening over a range of depths in the gas. Off-axis this broadening is worsened due to the contributions of the increased scattering and diffraction spikes from the mirror supports. The effective PSPC spatial resolution is a combination of (a) mirror scattering, (b) focusing and photon penetration effects, and (c) intrinsic broadening plus mirror blurring effects. These contribute various terms to the PSF with focussing and scattering coming to dominate
CHAPTER 3. X-RAY INSTRUMENTATION

Figure 3.4: ROSAT PSPC cross-section. The potentials of the grids are shown on the left and the grid thicknesses are shown on the right. The window composition includes a film of graphite to prevent electrostatic build-up. (Figure adapted from the ROSAT User's Handbook Chpt 3.)

beyond 14' of-axis. Approximately the PSF can be represented with a gaussian of FWHM 1' on-axis broadening to a few arcminutes off-axis (strongly depending on energy).

The PSPC has a FWHM spectral resolution, $\Delta E/E = 0.43 \times (E/0.93)^{1/2}$ with sensitivity over an effective range 0.1-2.4keV. Although the response of the XRT+PSPC extends beyond (see Fig 3.5) this range, calibration uncertainties make the data unusable for most cases. The low energy sensitivity is dominated by Carbon-K edge absorption in the entrance window and is complicated by some leak UV effects which can register in the lower energy channels. In order to prevent changes to the detector gas composition during operation from leakage or outgassing, the mixture is continually renewed.
from tanks at the rate of a few cm³/min.

Figure 3.5: PSPC spectral response. The low energy response is dominated by the C-K edge absorption feature. The high energy response of the detector cuts off rapidly after 2keV due to the diminishing effective area of the XRT. (Figure taken from the ROSAT User’s Handbook Chpt 3.)

3.4.4 ROSAT - High Resolution Imager

The second X-ray imaging system on ROSAT is the High Resolution Imager (HRI). It is an evolution of that flown on Einstein with the only significant change being the use of a CsI photocathode to improve quantum efficiency. The 'chevron' arrangement of micro-channel plates (MCP) is used to prevent feedback since the detector is operating at a very high gain of \( \sim 10^7 \).
charge is collected with a position sensitive multi wire grid in a similar way to the PSPC though the finer pitch of the grid allows finer spatial resolution to be achieved.

The effective spatial resolution of the HRI is affected by similar XRT blurring effects to the PSPC. The smaller diameter of the HRI detector plate itself means that the focal surface curvature is less important. Nevertheless, for sources further than 5' offaxis the PSF begins to increase noticeably above the on-axis PSF. The actual in-flight PSF and problems associated with source blurring are discussed in greater detail in Chapter 4.

Despite the low sensitivity of MCPs to changes in X-ray energy, some limited energy resolution is available. Fig 3.7 and Table 3.2 show the PHA distributions and mean PHA channel for various monochromatic sources determined from pre-flight calibration. The effective gain of the HRI gradually reduced during the mission due to degradation of the photocathode response leading to a shift of approximately 0.5 channel/year in the resulting PHA distribution (Prestwich et al. 1996). In June 1994, the HRI Gain Voltage was increased to compensate for the decline and this change restored the PHA distribution to that seen after launch.

The HRI instrument background for high and low rates of external particle events is shown in the lower two graphs in Fig 3.6. The background appears in general in the lower PHA channels since cosmic ray events and radioactive decay in the glass of the MCP can start cascades anywhere along the channel length and this in general will lead to lower total charge production than
### Table 3.2: The mean ROSAT HRI PHA channel for various monochromatic sources determined during ground calibration. The PHA distributions are shown in Fig 3.7 (from ROSAT HRI Spectral Calibration Report ftp://heasarc.gsfc.nasa.gov/rosat/software/hri/hrispec.tar.gz)

<table>
<thead>
<tr>
<th>Source</th>
<th>Energy</th>
<th>Mean PHA</th>
</tr>
</thead>
<tbody>
<tr>
<td>B</td>
<td>0.18keV</td>
<td>3.30</td>
</tr>
<tr>
<td>C</td>
<td>0.28keV</td>
<td>3.96</td>
</tr>
<tr>
<td>O</td>
<td>0.53keV</td>
<td>3.99</td>
</tr>
<tr>
<td>Fe</td>
<td>0.71keV</td>
<td>4.15</td>
</tr>
<tr>
<td>Cu</td>
<td>0.93keV</td>
<td>4.58</td>
</tr>
<tr>
<td>Al</td>
<td>1.49keV</td>
<td>5.44</td>
</tr>
<tr>
<td>Si</td>
<td>1.74keV</td>
<td>4.82</td>
</tr>
</tbody>
</table>

photon-triggered cascades from the top of a channel.

The spectral response of the detector varies across the surface and detailed gain maps have been made using the Bright Earth observations which are dominated by scattered solar X-rays at 525eV and provide a near-monochromatic flat field with which to monitor the instrument gain (Prestwich et al. 1996).

### 3.4.5 ASCA

The Japanese Advanced Satellite for Cosmology and Astrophysics (ASCA) was launched in February 1993 and operated until March 2001. It comprises four separate Wolter I XRT assemblies (focal length 3.5m) each with 120 foil mirror shells. Each of the XRTs focussing onto one of the detectors, two CCD imaging spectrometers (SIS0,SIS1) and two GSPCs (GIS2,GIS3) (Tanaka, Inoue & Holt 1994). Since the reflecting surfaces are thin foils, they
Figure 3.6: The ROSAT HRI background PHA distribution for high background (dashed line), low background (dotted line) and the source AR Lac (solid line) are shown as fractions of the total counts observed within that source. The source observation is far more strongly peaked than the two background distributions (from ROSAT HRI Spectral Calibration Report).
Figure 3.7: The ROSAT HRI PHA distributions measured during ground calibration from sources at various monochromatic energies. The dotted graph shown with boron (top left) is the PHA distribution from a 1.49keV aluminium source (from ROSAT HRI Spectral Calibration Report).
cannot be polished and as a result the surface roughness means the XRT PSF is approx 3' FWHM.

One major drawback with the ASCA XRT is that it is not equipped with pre-collimators which would absorb off-axis light before it entered the mirror nests. This means that stray light from bright sources outside the field of view can arrive at the focal plane after reflecting off just one set of mirrors. A single reflection will not bring light to a focus and instead increases the background across the detector.

![Figure 3.8: The effective area of the ASCA XRT at various off-axis angles. The sharp drop in effective area above 2keV is due to the M-edges of the gold reflective surface (from ASCA Technical Description Appendix E Fig 5.3a). The low energy cut-off for the GIS is mainly governed by the 10μm beryllium foil entrance window. This window still has a transmission of 10% down to](image-url)
0.7keV and is needed to seal the gas into the detector.

In the SIS, the main advantages are that it samples the XRT PSF without any further broadening and has better intrinsic spectral resolution than the GIS. The main disadvantage is that over the lifetime of the mission, $\Delta E$ has risen from 120eV at launch to $\sim$200eV.

### 3.4.6 Chandra

The Chandra satellite (previously known as AXAF) was launched in July 1999 and continues to operate. The most important advance over previous observatories is the very high spatial resolution of the High Resolution Mirror Assembly (HRMA). The Chandra HRMA consists of 4 nested Wolter type I mirror pairs with an iridium coating on glass. This achieves an image resolution of 0.5" FWHM on-axis.

The only detector referred to in this thesis is the Advanced CCD Imaging Spectrometer (ACIS) which has an energy range of 0.2-10keV and a spatial resolution of $\sim$2". There are two ACIS detectors on Chandra arranged in different configurations at the focal surface. Each is comprised of several separate CCD chips which are butted together. The ACIS-S set of CCDs are intended principally for high resolution spectroscopy with the transmission gratings in the beam path so rather than follow the primary focal surface of the HRMA, it follows the Rowland circle for the gratings. For on-axis imaging (on the S3 CCD), there is relatively little blurring from this arrangement.
The S3 CCD is at the on-axis position and is one of two backside illuminated CCDs which have been manufactured to leave the photo-sensitive depletion region exposed to the incoming X-rays. This provides better sensitivity to lower X-ray energies (see Fig 3.9). Due to damage from particles early in the Chandra mission, the front-side illuminated (FI) chips suffered a degradation in their energy resolution leading to some observers opting for the ACIS-S to make use of the undamaged back-side illuminated (BI) S3 chip. The effective energy resolution of the BI chips is $\Delta E/E \simeq 0.1$ at 1keV and 0.03 at 6keV.
In general, all of the above X-ray imagers produce a list of detected events giving position, arrival time and energy information for each. In most cases this list will have been pre-screened by the detector electronics (e.g. anti-coincidence, rise time) and subsequent software processing to remove events due to cosmic rays, $\gamma$ rays, UV photons or other 'undesirable' event triggers in the detector.

To detect source positions, the list of arrival coordinates is binned into an image array. The choice of bin size for this array is normally made to slightly oversample the PSF of the detector. Undersampling the PSF means that it is not possible to determine source extension to the limit of the instrument and excessive oversampling means there are too few events per bin (at the flux sensitivity limit of the detector) to provide sufficient statistics for detection.

In the simplest source detection approach, a 'sliding cell' is moved across the image and the counts, $(S + B)$ within the cell (of area $s$) are attributed to the source plus the background. Assuming that the background is constant over the image, those counts, $A$, from a source-free annulus (of area $b$) around the cell can be used to calculate the expected background, $\bar{B}$, within the source cell as $\bar{B} = A \cdot \frac{s}{b}$. The counts expected to be due to the source are then equal to the counts within the source cell minus the equivalent mean background for the cell area. Source significance is quantified in terms of the
signal-to-noise ratio,

\[ S/N = \frac{(S + B) - A \sigma^2}{\sqrt{(S + B) + A \sigma^2}} \]  

(3.9)

Where \((S + B)\) has more than 10 to 20 counts then the Gaussian approximation can be used (e.g. §6.3 Babu & Feigelson 1996). For fewer counts, Poisson statistics must be applied (Gehrels 1986). Both of these approaches still rely on classical statistics and more recent approaches based on Bayesian methods are improving the sensitivity of tests on low signal data (Nousek 1992, Marshall 1992).

Since the background is rarely smooth across the image, a more complete model of the background can be prepared using a smoothed map of the image with 'sliding cell'-detected sources removed. Local excesses in the counts can be compared with the smoothed background map to determine the probability of the source being real or a background fluctuation. In particular for the EXSAS software (Zimmermann et al. 1993) used for analysing the ROSAT data, Poisson-based Maximum Likelihood (ML) methods (Cash 1979, Crudace et al. 1988) are used. Here the probability, \(P\), that the measured photon distribution is due to a particular source strength is represented by a likelihood,

\[ L = -\ln(P) \]  

(3.10)

The Maximum Likelihood is defined as

\[ ML = L(S) - L(0) \]  

(3.11)
where $L(S)$ represents the likelihood of the distribution being observed for a source strength, $S$, and $L(0)$ being the likelihood of finding the distribution from the background strength. Extensive image simulation has led to probabilities being assigned to the measured ML in the EXSAS algorithms showing that ML values of 6, 10 and 14 correspond approximately to 3, 4 and 5 Gaussian sigma.

For the X-ray detectors discussed here, this situation is complicated by varying PSFs and backgrounds across the detector, making it difficult to provide a fully analytical description of the detection. In these cases extensive numerical simulation is used to calibrate the source detection procedures.

For the ASCA detectors, the presence of a large diffuse background component coupled with the varying gain across the detectors mean that the simple sliding cell approaches are not sufficient to determine the background. In particular, library backgrounds observed in source-free regions of the sky can be used. This does not help however when a diffuse source component or higher instrument background are present in an observation.

In the Chandra ACIS imager, the PSF is sufficiently small ($\sim 2''$) that much of the soft X-ray background is resolved into discrete sources. The remaining contribution from unresolved emission and instrumental background must still be subtracted.
3.6 X-ray Spectroscopy

Producing the spectrum of a source involves extracting the photons from a region around the source, and from a source-free region representing the background. These groups of photons are binned by energy channel and the difference between them gives the pulse height distribution of the source. Problems may arise as discussed in the previous section from incorrect background subtraction particularly in cases where there is non-uniformity or significant scattered light in the background (e.g. ASCA).

In order to decide if this distribution of counts is consistent with a particular spectrum, a model spectrum folded with the instrument energy response function is produced. This is compared with the source distribution and the model parameters adjusted until the best fit is achieved. The fit quality is usually determined from the reduced chi squared, $\chi^2/\nu$, where $\nu$ is the number of degrees of freedom calculated as the number of data bins minus the number of free spectral model parameters. In order to ensure that gaussian statistics can be used, each spectral bin should contain at least 20 counts. On the other hand if spectral bins are made too large, narrow spectral features may not be resolved (or may be lost entirely). Lampton, Margon & Bowyer (1976) Section IV.b conclude from simulations that since each energy channel is an independent sample of the random spectrum that increasing number of spectral bins still represent independent degrees of freedom. Similar conclusions are reached in simulations of ROSAT PSPC spectral fitting (Maggio et al. 1995).
Nevertheless, overbinning will result in increased counting errors per bin and loosen the constraints for spectral fitting. The optimal binning spectral bin size for a Gaussian redistribution is $\sim \Delta E / 3$ (Kaastra 1999). The typical bin sizes for the detectors described in $\S 2.3$ are shown in Table 2.3. Increasing the number of bins, $n$, beyond the optimal value will lead to a reduction in $\chi^2 / \nu$ by increasing $\nu$ and $\sigma$ per bin (thus reducing $\chi^2$) and giving a misleading goodness-of-fit.

For low energy resolution PSPC detectors this places a strong limit on the number of available bins for spectral fitting and hence on the complexity of spectral model that can be usefully constrained.

In practice, for the higher spectral resolution detectors, the rebinning is largely dominated by the available counts and the need to have sufficient counts per bin to justify the use of Gaussian statistics in spectral fitting. In this case energy channels are grouped together until a minimum number of counts per bin is achieved.

For faint sources where there are inadequate counts to produce a useful spectrum, hardness ratios can be calculated by comparing the detected counts in two wavelength bands. For the soft X-ray band this hardness ratio is
typically taken as,

\[ HR = \frac{H - S}{H + S} \]  \hspace{1cm} (3.12)

where H and S represent the detected counts in the soft and hard bands respectively.

### 3.7 Comparability of X-ray measurements made with different detectors

The widely differing energy ranges, sensitivities, PSFs, background characteristics and energy resolutions of the detectors introduce a number of problems when comparing fluxes from different instruments and observatories. In principle, the flux determined for each instrument has been corrected with response matrices and should represent the radiation that entered the front of the telescope. In practice, some factors will not have been taken into account and will mean that two identical fluxes will be 'seen' differently by two different detectors. Comparing results from the same detector removes most of the systematic variations though changes in response over the mission lifetime also affect comparison.

There are a number of issues with the low energy calibration of the ROSAT PSPC which complicate spectral modelling and absolute flux determination. The original spectral response of the PSPC was measured prior to launch at four energies at the PANTER facility of MPE in Garching and the remaining response model was an interpolation of these measurements. Much of the
calibration uncertainty in flight results from unknown contributions from scattered solar and auroral X-rays in the low energy channels. There is some disagreement on the comparability of spectral results taken by the ROSAT PSPC and ASCA SIS from simultaneous observations of NGC 5548 (Iwasawa, Fabian & Nandra 1999) but this may be due to specific spectral features being emphasised differently when folded with the two instrument responses (Miyaji, Lehmann & Hasinger 2001). However, good joint fitting result for the ROSAT PSPC and GINGA LAC (Large Area Proportional Counter) have also been presented (Pounds et al. 1994).
Chapter 4

ROSAT HRI Resolution
Recovery Techniques

4.1 Introduction

The ROSAT spacecraft was designed to observe sources while oscillating its pointing direction in one plane over a period of \( \sim 400 \text{ seconds} \) through \( \pm 1.5' \) (HRI) and by \( \pm 3' \) (PSPC). This was to prevent shadowing of sources behind the window supports in the PSPC and to prevent MCP pore burn-in for the HRI (as well as use more parts of the detector plate to limit the effects of local gain anomalies.) This observing mode means that source positions are recorded across a strip of the detector.

The spacecraft attitude is determined using the two startracker cameras (STCs) which imaged starfields 6° off the main spacecraft (and XRT) axis.
The defocussed star images were imaged across several of the 1' x 1' STC pixels and the position calculated from the centroid of the readings from each pixel. No image was record from the STCs with just the calculated star positions being logged every 1 second of observing time. Errors in this position determination cannot be corrected for using any of the other spacecraft engineering telemetry. If less than three guide stars were acquired by the startrackers then the data was not used for scientific analysis. The resulting spacecraft attitude log was used to calculate the sky position of each photon based on its arrival time and these processed attitude logs and photon events data were then included in the public release data.

From ground calibration, the ROSAT High Resolution Imager (HRI) was expected to have an on-axis PSF characterised as,

$$\sigma_{total}^2 = \sigma_{HRI}^2 + \sigma_{XRT}^2 + \sigma_{ASP}^2,$$

where, $\sigma_{HRI}=0.74''$, $\sigma_{XRT}=1.3''$ and $\sigma_{ASP}=1.5''$ are due to the HRI detector, the ROSAT XRT PSF and the spacecraft aspect error respectively. These combined to give an expected PSF described by a radial gaussian with $\sigma_{total}=2.12''$ (ROSAT Call for Proposals Appendix F). In practice the HRI rarely achieved actually source PSFs of the nominal size.

The initial in-orbit calibration using observations of the bright (12cts/sec) DA white dwarf HZ43 showed some broadening of the PSF beyond the expected ground calibration model of a point source and this was attributed to spacecraft aspect solution uncertainties. These observations were used
to adjust the effective model PSF used in the various data analysis suites (SASS, FTOOLS, EXSAS, etc) (David et al. 1997).

Most scientific observations were of far weaker sources (<1ct/sec) and taken over longer integration times. Many observations showed source broadening and asymmetries over and above what was expected from the calibrations. These were ascribed to spacecraft aspect solution effects due to incorrect tracking of guide stars through the 1’ pixels of the startrackers (Appendix A, Worrall et al. 1999, G. Hasinger, priv. comm.).

In 1998 it was realised that some of the problems were due to a bug in the SASS software which led to photon positions being systematically offset. As it happens the long calibration observations were taken in 1990 and 1992, periods when the effects of the software bug were minimised due to the small value of the fractional part of the spacecraft time (0.15625s and 0.140625s for 1990 and 1992 respectively, see Table 4.1).

In the following sections, I describe the three sources of image broadening for the ROSAT HRI and the software written here to analyse and correct these problems. Applications of the procedures to some bright test sources are discussed.

4.2 Aspect time error

This error in the calculated photon positions is due to a SASS software error in the processing of the original detector coordinates into Right Ascension
## CHAPTER 4. ROSAT HRI RESOLUTION RECOVERY

<table>
<thead>
<tr>
<th>Reset of S/C clock</th>
<th>Good Aspect Time</th>
<th>Bad Aspect Time</th>
<th>Difference</th>
</tr>
</thead>
<tbody>
<tr>
<td>90 Jun 01 (Launch)</td>
<td>0.15625</td>
<td>0.00122</td>
<td>0.155</td>
</tr>
<tr>
<td>91 Jan 25.386331</td>
<td>0.6875</td>
<td>0.00537</td>
<td>0.682</td>
</tr>
<tr>
<td>92 Feb 11.353305</td>
<td>0.140625</td>
<td>0.00140</td>
<td>0.139</td>
</tr>
<tr>
<td>93 Jan 18.705978</td>
<td>0.9375</td>
<td>0.00732</td>
<td>0.930</td>
</tr>
<tr>
<td>94 Jan 19.631352</td>
<td>0.625</td>
<td>0.00488</td>
<td>0.620</td>
</tr>
<tr>
<td>95 Jan 18.169322</td>
<td>0.6875</td>
<td>0.00537</td>
<td>0.682</td>
</tr>
<tr>
<td>96 Jan 28.489871</td>
<td>0.609375</td>
<td>0.00476</td>
<td>0.605</td>
</tr>
<tr>
<td>97 Jan 16.069990</td>
<td>0</td>
<td>0</td>
<td>0</td>
</tr>
<tr>
<td>98 Jan 19.445738</td>
<td>0.046875</td>
<td>0.00037</td>
<td>0.046</td>
</tr>
<tr>
<td>98 Sep 20 HRI destroyed due to the Sun entering the field of view during an Attitude Control failure.</td>
<td></td>
<td></td>
<td></td>
</tr>
</tbody>
</table>

Table 4.1: The time errors arising from the SASS Aspect time bug and the expected PSF broadening to result from a perfect wobble pattern. The actual broadening will be larger due to the deviations from the nominal wobble and from photons having their positions interpolated between the wrong time entries due to the aspect time offset. (Adapted from Mossman 1999).

and declination sky coordinates. The bug remained undetected until after the last XRT startracker failed in April 1998. Attempts were made to use the ROSAT Wide Field Camera startracker for subsequent HRI observations and systematic differences were found between the WFC and XRT startracker in the calculated pointing for previous archived observations. This ultimately led to the SASS bug being noticed at SAO. However, this error remains in photon positions (*.bas.fits event lists) in the ROSAT Data Archive (H. Meuller (Astr. Inst. Potsdam) & M. Corcoran (Goddard) priv. comm.) though the archived aspect times (in *.anc.fits) are correct.

The effect of the bug is dependent on the fractional part of the spacecraft clock time which remained approximately constant apart from clock resets
at the beginning of each year. The time errors shown in Table 4.1 were intro-
troduced during the determination of the spacecraft aspect log during each
observation. The spacecraft pointing was recorded every second with a co-
ordinated and time stamp. The time stamp comprised an integer number
of seconds and a fractional part stored as an integer number of 1/64ths of a
second. During processing of these data by SASS, the time was calculated
as: Time = Int + Frac/8192, instead of, Time = Int + Frac/64. The net
result of this was that the calculated fractional part was close to zero for all
years.

Photon arrival times are interpolated between the relevant entries in this
spacecraft aspect table and since the spacecraft is slewing all the time as it is
wobbled through ±1.5° over a period of 402s (see Fig 4.1), the aspect log error
means that the photon arrival time is interpolated incorrectly. Depending on
the photon arrival time relative to the fractional part of the spacecraft time,
the interpolation will be (a) between the correct aspect table entries but by
the wrong amount (always), and possibly also (b) interpolated between the
wrong table entries. These two options generate offsets in opposite directions
but both are proportional to the slewing velocity (see Fig 4.2). The photon
offsets are in the opposite sense during the second half of the wobble phase.
As can be seen in Fig 4.3, there is a periodic and a random component to
the event offsets.

As a result of this error the basic photon event tables included in the archives
have the incorrect values of the spacecraft pointing calculated by SASS as-
associated with each event. The actual spacecraft time stored in the ancilliary
Figure 4.1: The typical triangular ROSAT 402sec wobble pattern of ±1.5'. The figures show the variation of spacecraft RA (left) and Dec (right) against wobble phase (arbitrary offset). Each point represents the spacecraft pointing at each second of the observations. The 402s period of the wobble is maintained throughout successive observation intervals (here totalling 18ksec livetime taken over 39hrs.) The parts of the data deviating markedly from the wobble pattern are from during guide star acquisition at the beginning of an OBI and are not included in the good time intervals for data analysis.
Figure 4.2: A histogram of the wobble speeds for the data shown in Fig 4.1. The spacecraft spends most of its time slewing at between 1"/sec and 2.5"/sec during the observation.

data files (*.anc.fits) have the correct spacecraft time as they were processed with a different part of the SASS system.

Scripts to correct this bug have been written by Frank Primini (SAO) and described in Mossman (1999). The procedure followed is to calculate the incorrect aspect times from the correct ones included in the ancilliary tables using the known spacecraft clock fractional time for the epoch of the observation. Each photon arrival time is then used to interpolate between the correct and incorrect aspect table entries to determine the correction needed. This is applied for each photon in the table which can then be used in the normal manner.

This procedure is reliable so long as the satellite roll angle is not changing
Figure 4.3: The distribution of photon coordinate corrections (in units of 0.5") showing a periodic component in phase with the wobble and a random broadening component. The striping is due to events being corrected to the nearest 0.5" pixel. These corrections are those applied to the data in Fig 4.1 by the code provided by F. Primini.

rapidly since the interpolation algorithm assumes that the velocity vector between one aspect table entry and the next has a constant direction.

4.3 OBI Guide star re-acquisition error.

At the beginning of each Observing Interval (OBI), the guide stars are acquired in the startracker cameras. These calculated positions from the STCs are used to adjust the spacecraft boresight until it has the correct pointing. From ground calibration this boresight position was expected to be accurate to a few arcseconds ($\sigma_{ASP}$ in Eqn 4.1).
During in-flight calibration however it was realised that even with the STCs indicating the spacecraft boresight was in nominal pointing direction, the detected image positions could be displaced by up to 10". Since the wobble pattern for that OBI will be relative to the initially acquired position, the whole OBI will be observed with the initial error in the pointing. For bright sources this is straightforward to detect since a clear source detection can be made for each OBI (see Table 4.2). Photon events can be corrected for these offsets to bring each detection to a common position (e.g. centroid). This improves the relative scatter among the OBIs (and hence the PSF), but since the absolute spacecraft pointing information has been degraded by the original error, the final source positions may still have offsets of several arcseconds from their true positions on the sky.

4.4 Startracker wobble error

Similar to the OBI re-acquisition error, the STC tracking of the 402second wobble (driven by the spacecraft momentum wheels) produces inaccurately calculated guide star centroids. If the spacecraft wobble is uniform (Fig 4.4) the guide star images should travel over the same pixels of the STC every 402s and so the systematic boresight errors should be the same for a given wobble phase. These problems were investigated initially by Morse (1994) and subsequently by Harris (1998) who showed that a significant fraction of the aspect error was in phase with the wobble of the satellite. At this stage the software aspect time bug was not known but that effect is normally
of smaller amplitude than the full in-phase error. The solution originally

proposed by Morse (1994) and (now available for various analysis software) is to split the observed photons according to wobble phase into a number of bins and carry out source detections with each group of events. If the offsets are statistically significant then each event subset can be recentred accordingly. This works well for bright sources (as in the case of OBI re-centring) since there are sufficient source counts per sub-image that the positional errors on the detections are small compared to the measured offsets. Fig 4.5 shows the detected y coordinate of HZ 43 with the 1418173h observation split into 16 phase bins with the offsets showing a systematic single period variation approximately in-phase with the spacecraft wobble.
Figure 4.5: The y coordinate of wobble pattern (black, identical to Fig 4.4), and the y coordinate (red) of HZ43 as detected in each of the 16 phase bins (red) in 141873h. The scale for the latter is magnified by 50 with the data showing a half amplitude of ~1.8sky_pixels and an error per bin of ~0.1sky_pixels.

4.5 Other approaches from the literature

Most authors who have acknowledged the aspect residual problem have either indicated it as a possible explanation for small scale extension or assymetry or have smoothed their data (typically with a Gaussian and \( \sigma \) anywhere from 1'' - 10'') to take into account the (probable but undetermined) larger effective PSF of the observation.

Apart from dewobbling using phase binning and the recentroiding of OBIs using the Morse and Harris prescriptions, a number of approaches have been
described in the literature to quantify the level of aspect residuals and/or restore images.

Worrall et al. 1999 and Canosa et al. (1999) fitted a grid of $\beta$-models ($\beta = 0.5, 0.667, 0.9, 1.1; r_0 = 0" - 10"$) to subsets of the data which have the same roll angle and were taken less than 3 days apart for analysis of HRI images of radio galaxies. They note that sources broadened by aspect errors can appear like $\beta$-models with $r_0$ in the range 5"-10" depending on the value of $\beta$ chosen.

No authors have presented successful methods for detecting and restoring OBI and wobble residuals for sources fainter than 0.1cts/sec or for strongly extended sources.

4.6 Restoring data from faint and extended sources

As sources get fainter, positional errors increase (due to counting statistics) and the statistical significance of a detected sub-image offset reduces. If wider bins are chosen then significance may improve but then when the photon events within the bin are corrected by the mean bin offset, some will be corrected by too much and some by too little. This will introduce a blurring term into the correction which reduces the effectiveness of the re-centring.

Due to operational constraints, OBIs rarely exceed a few thousand seconds
Figure 4.6: The HRI image (5" pixels) of NGC 4552 (600491h1) discussed in Chapter 5 showing extension around the central source even after correction for the aspect time error. The plot (right) shows the scatter in source positions detected within each individual observation interval. The 'pix' units in the plots are 0.5" and all the OBI detected positions are within the central four pixels of the image.

in duration so for sources fainter than 0.1cts/sec, there will only be a few hundred source photons to determine the position. For sources with less than 100 counts, the positional errors can exceed the offsets between OBIs (Harris et al. 1998, Harris 1999). This situation can be improved by stacking fainter sources to improve the signal quality in the peak if there has been little change in spacecraft roll angle during the observation. However sources with off-axis angles larger than 5' show noticeable PSF broadening and their use is of questionable value as they increase the background. It is also possible to select a subset of the total OBIs or live exposure time within one OBI if some part of the observation is particularly degraded.

For extended sources there is also a drop in positional accuracy due to the
lower contrast with the background. When an extended source is further broadened by OBI and wobble residuals the intrinsic source extension can become very hard to determine. The extension in the source shown in Fig 4.6 is partly intrinsic to the source and partly due to the offsets shown in the OBI positions. Nevertheless, a strong nuclear peak in the image allows peak positions to be determined. Some of the OBI detections with large errors do not have sufficient counts to give a reliable source position.

If other point sources are available in the image then they may be preferrable for use in calculating the positions unless they are so faint that the errors are larger than for the extended source.

4.7 Two 'half phase' bins

In order to determine whether OBI and wobble residual detection can be extended to sources much fainter than the 0.1cts/sec at which the publically available software routines are effective, a set of EXSAS/MIDAS routines were written to split the data into OBIs and wobble phase bins and examined the behaviour of these codes on the bright calibration source HZ43 (observations 141873h and 142549h).

For a faint source, the smallest binning that can be used is to split the data in two. This preserves the maximum source counts in each bin while still allowing some exploration of the wobble phase dependence of the source photon positions. Since the underlying wobble errors have an unknown phase
relationship with the spacecraft wobble phase, an arbitrary bin phase offset relative to the wobble phase, $\phi$, is chosen. All events with phase in the range $[\phi, \phi + 1/2]$ go in one bin and those in $[\phi - 1/2, \phi]$ go in the other.

Figure 4.7: The $y$ coordinates of HZ43 detected in each of two 'half wobble' phase bins as a function of the offset phase ($\%$) of the start of the bin. Positional errors are 0.1 sky-pix for each bin.

For the same data as shown in Fig 4.5, the detected coordinates in the two half phase bin sub-images change with $\delta$ with maximum difference being reached when each half phase bin contains the most separated parts of the wobble residuals. Since these residuals show the maximum and minimum differences being separated by a phase of $\sim 0.25$, the simplest initial working assumption was that the wobble residual is a triangular wave at some phase offset from the spacecraft wobble. In Fig 4.5 the actual detected phase offsets for 16
CHAPTER 4. ROSAT HRI RESOLUTION RECOVERY

<table>
<thead>
<tr>
<th>Bins</th>
<th>$y_{max}$</th>
<th>$\bar{y}$</th>
<th>$y_{min}$</th>
<th>$\phi_{max}$</th>
<th>$A$</th>
</tr>
</thead>
<tbody>
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<td>2</td>
<td>6.65</td>
<td>5.98</td>
<td>5.30</td>
<td>26.0%</td>
<td>2.70</td>
</tr>
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<td>4</td>
<td>6.74</td>
<td>5.96</td>
<td>5.17</td>
<td>25.0%</td>
<td>2.04</td>
</tr>
<tr>
<td>6</td>
<td>7.73</td>
<td>6.09</td>
<td>4.45</td>
<td>25.0%</td>
<td>3.94</td>
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<tr>
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<td>6.45</td>
<td>5.26</td>
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<td>8.07</td>
<td>6.43</td>
<td>4.78</td>
<td>29.1%</td>
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<tr>
<td>16</td>
<td>7.87</td>
<td>6.35</td>
<td>4.83</td>
<td>21.1%</td>
<td>1.62</td>
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</table>

Table 4.2: Phase dependence of offsets for different number of bins for HZ43. The amplitude, $A$, of the triangular residual is calculated as Eqn 4.2

Bins are clearly more complex than just a triangular profile but nevertheless it shows its maxima and minima around phases 0.25 and 0.75.

For a triangular profile, the amplitude of the residual can be calculated as,

$$A_n = A \left(1 + \frac{1}{n}\right) \quad (4.2)$$

where $n$ is an even number of bins and $A$ is the full peak-to-peak amplitude of the triangular wave.

This approach was tested on the 141873h data for 2, 4, 6, 8, 10, 12, 14 and 16 bins and the phase and amplitude of the maximum peak-to-peak distances measured in each case. The calculated values of the triangular residual model amplitude, $A$, are shown in Table 4.2. The 2 bin determinations show an overestimate of the amplitude compared to the 16 bin measurement although the phase at which maximum peak-to-peak difference is detected remains approximately constant.
4.8 Correcting a single OBI

Determining the wobble residuals for a whole observation (i.e. several OBIs) has the advantage of maximising the statistics available but the disadvantage of allowing OBI offsets to affect the source distribution. Fig 4.8 shows the declination component of the spacecraft wobble for the 141873h observation. The five OBIs have different wobble properties with only the last two showing full normal wobble behaviour.

Due to the high count rate of HZ43 it is possible to apply the dewobbling approach to a single OBI and still have enough counts to detect source positions accurately within several phase bins. In Fig 4.9 the data for HZ43 observation 142549h show that during OBI #1 the roll angle is approximately constant and a well defined wobble pattern is present for ~7.5 wobble periods.

Source detection parameters in each OBI are shown in Table 4.3 illustrating the clear displacement between the detected positions of the source between the two OBIs. There is a modest improvement in extension after applying the aspect time correction but most of this is due to the broader OBI #2 and the offsets between the two OBIs.

The wobble residual detected over 16 bins and the half phase binning for this data are shown in Fig 4.10. Again the detected maximum phase agree between the two methods (~27%) and the predicted triangular amplitudes are 3.1 and 4.2 respectively.

This 'half phase' bin detection method for determining the level of wobble
Table 4.3: The source detection parameters for the two 142549h OBIs. For illustration, both the raw and aspect time corrected (Primini script) data are used. The source detected position, $X_{\text{sky}}$, $Y_{\text{sky}}$, the position error, $\Delta X$, and the source extension, Ext, are all in units of $\text{sky.pix}$ ($0.5''$). The maximum likelihood for the extension EXT_ML shows clearly that the intrinsic extension is much lower for OBI #1 than #2.

<table>
<thead>
<tr>
<th></th>
<th>$X_{\text{sky}}$</th>
<th>$Y_{\text{sky}}$</th>
<th>$\Delta X$</th>
<th>Ext</th>
<th>EXT_ML</th>
<th>CTS</th>
</tr>
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<tbody>
<tr>
<td>Full</td>
<td>Raw</td>
<td>120.26</td>
<td>10.73</td>
<td>0.1</td>
<td>0.70±0.11</td>
<td>2951.5</td>
</tr>
<tr>
<td></td>
<td>Primini</td>
<td>120.14</td>
<td>10.66</td>
<td>0.1</td>
<td>0.70±0.11</td>
<td>2884.3</td>
</tr>
<tr>
<td>OBI#1</td>
<td>Raw</td>
<td>121.88</td>
<td>11.32</td>
<td>0.2</td>
<td>0.57±0.13</td>
<td>467.8</td>
</tr>
<tr>
<td></td>
<td>Primini</td>
<td>121.90</td>
<td>10.73</td>
<td>0.2</td>
<td>0.50±0.12</td>
<td>328.7</td>
</tr>
<tr>
<td>OBI#2</td>
<td>Raw</td>
<td>119.39</td>
<td>11.00</td>
<td>0.2</td>
<td>0.77±0.13</td>
<td>2626.1</td>
</tr>
<tr>
<td></td>
<td>Primini</td>
<td>119.45</td>
<td>10.66</td>
<td>0.2</td>
<td>0.74±0.12</td>
<td>2336.0</td>
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</table>

residuals allows the approach to be extended to much fainter sources since the original source counts need only be halved in each sub-image rather the standard approaches of 5, 10 or 20 bins.
Figure 4.8: The spacecraft declination as a function of time and wobble phase for the 141873p HZ43 observation. Each point represents the spacecraft position at one second intervals. The top graph shows the full timeline for the observation with the five OBIs. An enlarged view of the last two OBIs (bottom left) shows the normal satellite wobble behaviour. The declination as a function of phase (bottom right) for the 402second wobble period shows some of the observation un-wobbled and some wobbled.
Figure 4.9: The spacecraft roll angle (left) and declination (right) during OBI#1 of the 142549h observation.
Figure 4.10: The source positions (top) determined for OBI#1 of the 142549h observation in 16 phase bins (errors of 0.4sky_pix for each bin.) The approximate triangular shape of the residuals is clear. This same signature is present in the coordinates detected in each of the half phase bins (bottom).
Chapter 5

The soft X-ray variability of NGC 4552

5.1 Introduction

The giant elliptical galaxy, NGC 4552, is the tenth most luminous E/EO object in the Virgo cluster. Optically it shows no obvious signs of nuclear activity and has a spectrum dominated by broad absorption lines typical of old stellar populations in ellipticals (Davies et al. 1987, Ho, Filippenko & Sargent 1995). The central ~2.5kpc shows significant Hα emission from the ISM (Macchetto et al. 1996). The galaxy also has a weak compact radio source which has shown variability on a timescale of years (Jenkins 1982, Wrobel 1991, Wrobel & Heeschen 1991). More recently, observations of this variable core have also revealed evidence of parsec-scale jets (Nagar et al.
2000, Nagar, Wilson & Falcke 2001, Nagar et al. 2002). At X-ray energies, a hot ISM component typical of ellipticals was detected by Einstein (Forman, Jones & Tucker 1985) at a best fit temperature of 0.9keV (Raymond-Smith) (Kim, Fabbiano & Trinchieri 1992).

Figure 5.1: The discovery images taken on 19th July, 1991 (left) and 28th November, 1993 (right) by the HST/FOC of the varying point source at the centre of NGC 4552 from Renzini et al. (1995). Each image is 2.9" square with the central peaks coinciding with the centroid of the galaxy isophotes to within one pixel (0.02"). The off-centre source in the 1991 image is 0.14" from the centre. North is up and East is left in both images.

In 1993 NGC 4552 was observed to have a sharp 'spike' within ~2pc from its centre in UV images (Fig 5.1) taken by the Faint Object Camera (FOC) on the Hubble Space Telescope (HST) (Renzini et al. 1995). In comparison with a similar FOC image taken in 1991, this feature had brightened ~4.5 times. Follow-up observations in 1996 with the post-COSPAR FOC and Faint Object Spectrograph (FOS) showed a drop by a factor of 2 from the
1993 level (Cappellari et al. 1999). A second spike was also detected in the 1991 observation at a distance of 0.14" (~10pc) from the central spike. This was only detected in the F342W (U band) filter and was not detected in either the 1993 or 1996 follow up observations. The fluxes for the central spike are shown in Table 5.1.

<table>
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<th>Filter</th>
<th>$\lambda_0$ (Å)</th>
<th>FWHM (Å)</th>
<th>$\lambda_{eff}$ (Å)</th>
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<th>$f_{1993}$ 28.11.1993</th>
<th>$f_{1996}$ 23.06.1996</th>
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<td>594</td>
<td>2800</td>
<td>-</td>
<td>17.7±1.3</td>
<td>8.2±0.4</td>
</tr>
<tr>
<td>F342W</td>
<td>3410</td>
<td>702</td>
<td>3400</td>
<td>3.2±0.4</td>
<td>14.4±1.2</td>
<td>7.0±0.4</td>
</tr>
<tr>
<td>F502W</td>
<td>4940</td>
<td>530</td>
<td>4990</td>
<td>7.0±0.8</td>
<td>-</td>
<td>-</td>
</tr>
</tbody>
</table>

Table 5.1: Fluxes measured in the NGC 4552 spike in the three sets of FOC observations. The fluxes $f_{1991}$, $f_{1993}$ and $f_{1996}$ (in units of $10^{-18}$ erg.s$^{-1}$.cm$^{-2}$.Å$^{-1}$) are from the observations data from Cappellari et al. 1999 and FOC Handbook.

The 1996 FOS spectroscopy indicated that as well as narrow emission lines there were also broad emission lines. The line ratios of H$\alpha$ to [N II], [S II] and [O I] lead to a classification of the spike as an extreme AGN, and the [O III]/H$\beta$ ratio puts it on the border between Seyferts and LINERs. Fitting the galaxy-subtracted UV spectrum of the spike gives temperature of $\sim$15,000K which would imply a bolometric luminosity of $\sim$3.6x10$^6$L$_\odot$ at the Virgo cluster distance of 16.8Mpc (Tully 1988). The emission line profiles indicate broad (FWHM$\sim$3000km.s$^{-1}$) and narrow line components (FWHM$\sim$700km.s$^{-1}$) although the broad components are also present in the forbidden lines in contrast with the standard behaviour in an active nucleus.

These spectral features coupled with the flux variability suggest some form
of transient accretion as the underlying mechanism. Capellari et al. 1999 suggest that this could be provided by the tidal stripping of the atmosphere of a star as it passes outside the tidal radius, $r_T$ of the central black hole.

### 5.2 X-ray analysis

The stellar disruption models and candidates detected to date (Chapter 2) suggest that a soft excess could be expected in X-rays if the 1993 UV flare was due to stellar disruption. Specifically it would be interesting to see if there are any correlated variations in the UV and X-ray emission during the early 1990’s which might be associated with a disruption event. To check this, an analysis of the available X-ray archive observations of NGC 4552 was made.

NGC 4552 has been observed by several satellite observatories in the soft X-ray band since its first detection by Einstein in 1979. ROSAT has the longest sequence of soft X-ray observations of the galaxy with the RASS, 4xPSPC and 2xHRI pointed observations (See Fig 5.2 and Table 5.2). Results from the ROSAT data have been presented by Beuing et al. (1999) and Davis & White (1996) though in both cases rather simple spectral models were used. In the following sections I present the detailed analysis of the ROSAT X-ray data and correlate it with data from other observatories in terms of basic source parameters, spatial and spectral analysis and timing analysis where feasible. A discussion of the long term and UV-correlated variability and the overall implications of the observations are presented in the final section.
5.3 ROSAT All-Sky Survey

The ROSAT All-Sky Survey (Voges et al. 1999) was carried out between July 1990 and January 1991. The PSPC was used to observe the sky along great circles as the satellite roll axis was kept approximately parallel to the Earth-Sun line of centres. Over the six months of the survey the satellite observed the whole sky. At the ecliptic latitude of NGC 4552 (~15°) the sky was observed over three days for a total effective exposure time of ~400s. The exposure time map for the second processing of the All-Sky Survey (RASS II) is shown in Fig 5.3. The long exposure times (black) at the ecliptic poles where the survey scan strips overlap are clearly visible, as are the areas with zero or very short exposure time (white). The latter are due to automatic detector shutdown for passage through the South Atlantic Anomaly and auroral zones, and also where inadequate numbers of guide stars were available. The inset in Fig 5.3 shows the position of NGC 4552 on the edge of a zone of very low exposure. As a result no sources from this area are included in the RASS Bright Source Catalog (BSC) or Faint Source Catalog (FSC).

The final archive data release (1999) contained extra exposure time which filled some of the gaps in the previous data and this allowed an analysis of NGC 4552 to be carried out. The analysis procedures for properly dealing with the RASS calibration files were included in the April 2001 release of the EXSAS software (Zimmermann et al. 1992). Automatic exposure calculation (and therefore count rates) are not included in the standard tools so I have
Figure 5.2: The ROSAT All-Sky Survey image of the centre of the Virgo cluster showing the bright cluster emission around M87 (right of centre). The six ROSAT pointed observations of NGC 4552 (centre of frame), 4xPSPC and 2xHRI, are shown as 57' and 20' circles respectively.

<table>
<thead>
<tr>
<th>Start-End Dates</th>
<th>Pointing</th>
<th>Off Axis</th>
<th>Time</th>
</tr>
</thead>
<tbody>
<tr>
<td>1990.11.06-11.08</td>
<td>(RASS)</td>
<td>-</td>
<td>0.4ks</td>
</tr>
<tr>
<td>1991.12.15-12.16</td>
<td>700056p</td>
<td>51.98'</td>
<td>4.4ks</td>
</tr>
<tr>
<td>1992.07.06</td>
<td>600437p</td>
<td>40.02'</td>
<td>0.4ks</td>
</tr>
<tr>
<td>1992.12.19-12.21</td>
<td>600437p1</td>
<td>40.34'</td>
<td>18.0ks</td>
</tr>
<tr>
<td>1993.06.30-07.05</td>
<td>600586p</td>
<td>5.34'</td>
<td>16.7ks</td>
</tr>
<tr>
<td>1994.07.09-07.11</td>
<td>600491h</td>
<td>0.32'</td>
<td>18.2ks</td>
</tr>
<tr>
<td>1995.06.14-06.17</td>
<td>600491h1</td>
<td>0.29'</td>
<td>30.1ks</td>
</tr>
</tbody>
</table>

Table 5.2: Summary of the ROSAT observations of NGC 4552.
Figure 5.3: Exposure map plotted in equatorial coordinates (RA=0hrs at right and increasing to left) of the RASS2 data release used in creating the RASS-BSC and RASS-FSC. The greyscale colours shade from white (no exposure) to black (36ksec maximum), at the ecliptic poles. The 8.5° × 8.5° inset shows the gap in the exposure strip through the NGC 4552 position with the exposure contours marked in seconds.

described the procedures I have used for RASS data analysis throughout the thesis in Appendix A. The detailed analysis and results from the RASS data are presented and discussed with the PSPC pointed observations.
5.4 ROSAT PSPC Pointed Observations

5.4.1 Source detection

ROSAT carried out four pointed observations of NGC 4552 with the PSPC B detector. All were made between the July 1991 and November 1993 observations of the UV flare by HST/FOC with the final PSPC observation being made 150 days before the UV flare was observed at its maximum brightness.

The data for each observation were downloaded from the ROSAT Archive, and analysed using the EXSAS software (APR01 release). Detector response matrices PSPC C and PSPC B were used for the RASS and pointing data respectively. Binned images were created for photon events between channels 10 and 240 (0.1-2.4 keV) and the normal source detection pipelines were carried out with the DETECT/SOURCES maximum likelihood method.

NGC 4552 was detected in all full energy band images (0.1-2.4 keV) with $ML > 15$ and the source parameters are given in Table 5.3. The same source detection procedure was carried out for photon events extracted between 0.1-0.5 keV (Soft band) and 0.51-2.4 keV (Hard band) and Table 5.3 also lists these detection parameters.
<table>
<thead>
<tr>
<th></th>
<th>RASS</th>
<th>700056p</th>
<th>600437p</th>
<th>600437p1</th>
<th>600586p</th>
</tr>
</thead>
<tbody>
<tr>
<td><strong>Exposure</strong></td>
<td>443.7s</td>
<td>9313s</td>
<td>377s</td>
<td>17976s</td>
<td>16660s</td>
</tr>
<tr>
<td><strong>Offaxis</strong></td>
<td>-</td>
<td>52'</td>
<td>40.0'</td>
<td>40.3'</td>
<td>5.3'</td>
</tr>
<tr>
<td><strong>Livetime</strong></td>
<td>-</td>
<td>0.96855</td>
<td>0.97288</td>
<td>0.96992</td>
<td>0.95849</td>
</tr>
<tr>
<td><strong>Soft (0.1-0.5keV)</strong></td>
<td><strong>Counts</strong></td>
<td>&lt;11.19</td>
<td>&lt;44.79</td>
<td>&lt;18.36</td>
<td>&lt;68.06</td>
</tr>
<tr>
<td>ML</td>
<td>2.2</td>
<td>0.5</td>
<td>2.6</td>
<td>2.0</td>
<td>391.1</td>
</tr>
<tr>
<td>CR(/ksec)</td>
<td>&lt;25.2</td>
<td>&lt;8.4</td>
<td>&lt;75.9</td>
<td>&lt;6.1</td>
<td>23.0±1.4</td>
</tr>
<tr>
<td>VigCorr</td>
<td>1.517</td>
<td>1.692</td>
<td>1.522</td>
<td>1.555</td>
<td>1.018</td>
</tr>
<tr>
<td>Bgnd</td>
<td>2.419</td>
<td>0.962</td>
<td>1.021</td>
<td>1.092</td>
<td>1.564</td>
</tr>
<tr>
<td><strong>Hard (0.51-2.4keV)</strong></td>
<td><strong>Counts</strong></td>
<td>35.72±6.72</td>
<td>403.81±27.30</td>
<td>18.30±5.20</td>
<td>687.56±33.48</td>
</tr>
<tr>
<td>ML</td>
<td>53.8</td>
<td>191.4</td>
<td>14.7</td>
<td>420.7</td>
<td>4094.8</td>
</tr>
<tr>
<td>CR(/ksec)</td>
<td>80.5±15.1</td>
<td>96.3±6.5</td>
<td>81.1±23.1</td>
<td>64.0±3.1</td>
<td>83.5±2.4</td>
</tr>
<tr>
<td>VigCorr</td>
<td>1.592</td>
<td>2.151</td>
<td>1.632</td>
<td>1.627</td>
<td>1.021</td>
</tr>
<tr>
<td>Bgnd</td>
<td>1.399</td>
<td>0.415</td>
<td>0.646</td>
<td>0.584</td>
<td>0.782</td>
</tr>
<tr>
<td><strong>Full (0.1-2.4keV)</strong></td>
<td><strong>Counts</strong></td>
<td>38.95±7.84</td>
<td>229.36±31.26</td>
<td>28.48±7.01</td>
<td>708.95±41.77</td>
</tr>
<tr>
<td>ML</td>
<td>36.7</td>
<td>28.0</td>
<td>16.0</td>
<td>150.7</td>
<td>3890.1</td>
</tr>
<tr>
<td>CR(/ksec)</td>
<td>87.8±17.7</td>
<td>50.2±6.8(†)</td>
<td>119.3±29.4</td>
<td>63.0±3.7</td>
<td>106.4±2.7</td>
</tr>
<tr>
<td>VigCorr</td>
<td>1.556</td>
<td>1.974</td>
<td>1.542</td>
<td>1.555</td>
<td>1.020</td>
</tr>
<tr>
<td>Bgnd</td>
<td>3.799</td>
<td>1.574</td>
<td>1.601</td>
<td>1.667</td>
<td>2.345</td>
</tr>
</tbody>
</table>

Table 5.3: Source detections in each band for the ROSAT PSPC data. Results are presented for the soft (0.1-0.5keV), hard (0.51-2.4keV) and full (0.1-2.4keV) bands. Count rate is calculated as $\frac{Counts \times VigCorr}{Exposure \times Livetime}$ for the pointed exposures. For the RASS count rate is, as the exposure is calculated from the vignetting corrected exposure map. If there is no detection in the soft band at ML>10 then the 95% confidence upper limit to the counts is shown. The background count rate is shown in units of cts/ksec/arcmin$^2$. †The Full band CR is lower than the Soft + Hard sum due to the very low source count contribution from the soft band relative to the high soft background contribution, leading to a lower detected count rate in the full band.
5.4.2 Extension

For the first three PSPC pointings, large off-axis angles (and for 700056p and 600437p, short exposure times also) reduce the quality of the data. The image contours shown in Fig 5.4 show significant distortion of the 700056p, 600437p and 600437p1 images due to the off-axis PSPC PSF. The RASS and 600586p images show a clear core surrounded by fainter emission with some indications of North-South extension. Neither of the long exposures is consistent with a point source PSF with detected extensions (1σ) of 13.7±3.5” (600586p) and 59.1±24.6” (600437p1). The other observations have too few counts or are too far off-axis (and hence too distorted and spread over too large a background area) to detect extension at any level of significance.

In the 700056p observation, NGC 4552 is close to the edge of the detector and the vignetting changes across the source. The galaxy is close to a window support rib in the outer part of the detector in both 600437p and the follow up 600437p1 which also distorted the extension profile.

5.4.3 Variability

Light curves were made for NGC 4552 in each observation and were binned by the 400sec wobble period. No significant variability was seen over the duration of each observation. The light curve for observation taken over the longest timebase (600586p, ~4.6 days) is shown in Fig 5.5 and is consistent with a constant count rate.
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Figure 5.4: Source detection images (400" x 400") from the PSPC observations; RASS (top), 700056p (mid-left), 600437p (mid-right), 600437p1 (bottom left) and 600586p (bottom right). Contours are at 1.2, 1.5, 2, 3 and 5 times the background in each case.
CHAPTER 5. THE SOFT X-RAY VARIABILITY OF NGC 4552

Figure 5.5: The light curve 600586p in 400s bins. The upper data points show the binned count rates for the source extraction region before background removal. The lower points are the binned rates from the background extraction region.

To examine the possibility of systematic problems affecting the detected NGC 4552 count rates in the PSPC observations, other source detections in each of the observations were compared. Table 5.4 shows six sources detected above ML=15 in at least two observations and the detections or upper limits for non-detections if the source was in the FOV of an observation. The optical identifications (within 2’) from the NASA Extragalactic Database show three of the objects with active classifications. The count rates vary by a factor of up to ~2.8. There are some indications of systematic change in count rate between 700056p and 600586p with sources A, B and C showing evidence for brightening. Similarly sources A and E show the same trend between 600437p1 and 600586p although source F is consistent with no change. Sources A and G show opposite variations between 600437p and 600437p1.
### Table 5.4: Other detections in the PSPC frames. Count rates are shown as counts per ksec. For the two non-detections in 600437p 95% confidence upper limits are given. The optical identifications and activity types are taken from NED.

<table>
<thead>
<tr>
<th></th>
<th></th>
<th></th>
<th></th>
<th></th>
</tr>
</thead>
<tbody>
<tr>
<td>A</td>
<td>12h35m39.7s</td>
<td>12d33m27.9s</td>
<td>NGC 4552</td>
<td>E</td>
</tr>
<tr>
<td>B</td>
<td>12h37m43.6s</td>
<td>11d49m02.0s</td>
<td>M58/NGC 4579</td>
<td>LINER/Sy2</td>
</tr>
<tr>
<td>C</td>
<td>12h34m52.00s</td>
<td>11d57m24.9s</td>
<td>RMB123</td>
<td></td>
</tr>
<tr>
<td>D</td>
<td>12h37m01.8s</td>
<td>12d00m47.0s</td>
<td>LBQS1234+1217</td>
<td>QSO</td>
</tr>
<tr>
<td>E</td>
<td>12h36m49.58s</td>
<td>13d09m48.7s</td>
<td>M90/NGC 4569</td>
<td>LINER/Sy</td>
</tr>
<tr>
<td>F</td>
<td>12h35m36.17s</td>
<td>13d11m02.2s</td>
<td></td>
<td></td>
</tr>
<tr>
<td>G</td>
<td>12h35m58.58s</td>
<td>13d29m15.3s</td>
<td>VPC1028/VCC1650</td>
<td></td>
</tr>
<tr>
<td>A</td>
<td>700056p</td>
<td>600437p</td>
<td>600437p1</td>
<td>600586p</td>
</tr>
<tr>
<td>B</td>
<td>50.2±6.8</td>
<td>119.3±29.4</td>
<td>63.0±3.7</td>
<td>106.4±2.7</td>
</tr>
<tr>
<td>C</td>
<td>378±7</td>
<td>623±8</td>
<td></td>
<td></td>
</tr>
<tr>
<td>D</td>
<td>9±2</td>
<td>24±4</td>
<td></td>
<td></td>
</tr>
<tr>
<td>E</td>
<td>9±2</td>
<td>12±1</td>
<td></td>
<td></td>
</tr>
<tr>
<td>F</td>
<td>&lt;33</td>
<td>29±2</td>
<td>75±5</td>
<td></td>
</tr>
<tr>
<td>G</td>
<td>&lt;12</td>
<td>10±1</td>
<td>11±2</td>
<td></td>
</tr>
<tr>
<td></td>
<td>100±20</td>
<td>267±4</td>
<td></td>
<td></td>
</tr>
</tbody>
</table>

The small number of events detected for the 600437p and 600437p1 observations mean that no useful spectral content was obtained. The statistical significance of the events were estimated using $S/N > 3$ criteria. The PSPC count rates were calculated on a $10^{-4}$ level, where if and $i$ are the counts detected in the bands and $i$ is the background rate.
5.4.4 Spectral analysis

The small numbers of counts detected for the RASS, 700056p and 600437p observations mean that no useful spectra can be fitted. The usual statistical binning with S/N>5 resulted in only 2 to 4 bins. Instead hardness ratios are calculated as \( \frac{H-S}{H+S} \), where H and S are the counts detected in the hard and soft bands. An equivalent ratio is also calculated to allow comparison where there is no soft band detection using the ratio \( \frac{H-(F-H)}{F} \). The two ratios should give similar values when sufficient counts are available for reliable source detection.

The 600437p1 observation has a greater number of counts but again partly due to the fact that it is spread over a wider area, very poor constraints were placed on the models fitted and the hardness ratio method is used for this data also.

<table>
<thead>
<tr>
<th></th>
<th>RASS</th>
<th>700056p</th>
<th>600437p</th>
<th>600437p1</th>
<th>600586p</th>
</tr>
</thead>
<tbody>
<tr>
<td>HR</td>
<td>( \frac{H-S}{H+S} )</td>
<td>&gt;0.52</td>
<td>&gt;0.80</td>
<td>&gt;0.00</td>
<td>&gt;0.82</td>
</tr>
<tr>
<td></td>
<td>( \frac{H-(F-H)}{F} )</td>
<td>0.83±0.49</td>
<td>Undefined</td>
<td>0.29±0.22</td>
<td>0.94±0.16</td>
</tr>
</tbody>
</table>

Table 5.5: Hardness Ratios for PSPC detections

The 600586p observation has 1670 counts in the range 0.1-2.4keV and source events were extracted from a circular region of radius 120" around the detected source position. This large value was used to ensure that all the extended emission was included in the extraction cell. Background events were extracted in an annulus of radius 300" around the source cell.

The initial spectral binning chosen was the standard S/N ratio of 5. This
gave 49 bins with the strongly peaked region around 1 keV having bin sizes of 0.2 keV. S/N ratios of 8 and 9 were also used for binning and produced broader bin widths. Nevertheless, the S/N = 9 distribution still had narrow bins around 1 keV so these were combined with the neighboring bin to get a minimum binsize of 0.1 keV. This produced a distribution with 13 bins.

Fits to the data were made for a single absorbed powerlaw (PL), a single absorbed Raymond-Smith (RS) thermal plasma and a two component absorbed (PL+RS) model. The absorption cross-sections used are from Morrison & McCammon (1983). The results for fitting these models are shown in Table 5.5. All parameters were free during the fitting but no convergence was achieved if the abundance was free so this was fixed to solar abundance and the remaining parameters fitted. The best fit to the two component model is shown in Fig 5.6.

Davis & White (1996) have presented an analysis of the 600586p observation in which they modelled the spectrum with a single absorbed Raymond-Smith thermal plasma \( (kT=0.74^{+0.07}_{-0.06} \text{ keV}, N_H=3.10^{+0.77}_{-0.80} \times 10^{20} \text{ cm}^{-2} \) with abundance of \( 0.08^{+0.07}_{-0.03} \) solar). It should be noted that they have used 51 degrees of freedom in their spectral fitting.\(^1\)

\(^1\)Following from Chpt 3.6, the optimal binsizes are 0.20 keV, 0.14 keV and 0.11 keV for the ROSAT PSPC (FWHM \( \Delta E/E \approx 0.43 \left( \frac{E}{0.33} \right)^{-0.5} \) at energies 0.5, 1.0 and 1.5 keV respectively. This means that there should be no more than \( \sim 20 \) bins for a PSPC spectrum even if signal to noise considerations permit higher binning as is the case here.
Figure 5.6: Power law plus Raymond Smith models with absorption fitted to the 600586p spectrum. The dashed line shows the 0.70keV Raymond Smith plasma component and the dotted line shows the $\Gamma=2.36$ power law component. The rebinning for this data was done with S/N=9-. 
Table 5.6: Spectral fitting parameters for the 600586p observation of NGC 4552. The relative strengths of the two spectral components are indicated by the normalisations of each model at 1keV. The PL normalisation is in units of $10^{-5}$ photons cm$^{-2}$ s$^{-1}$ keV$^{-1}$. The RS normalisation is in units of $\frac{10^{-18} \text{ cm}^{-5} \text{ s}^{-1}}{4\pi D_L^2} f_V n_e n_H dV$. The S/N ratio used for binning is indicated in brackets after the model name with S/N=5,8,9. S/N=9- is the manual rebinned S/N=9 distribution as described in the text.

<table>
<thead>
<tr>
<th>Model</th>
<th>$N_H$ ($10^{20}$ cm$^{-2}$)</th>
<th>$kT$ (keV)</th>
<th>RS Norm.</th>
<th>PL Norm.</th>
<th>$\chi^2/\nu (\nu)$</th>
</tr>
</thead>
<tbody>
<tr>
<td>RS(9-)</td>
<td>0.55±0.21</td>
<td>1.00±0.05</td>
<td>3.91±0.54</td>
<td>-</td>
<td>5.54±0.34</td>
</tr>
<tr>
<td>PL(9-)</td>
<td>7.56±1.06</td>
<td>-</td>
<td>-</td>
<td>2.92±0.19</td>
<td>5.54±0.34</td>
</tr>
<tr>
<td>RS+PL(5)</td>
<td>4.02±1.07</td>
<td>0.66±0.33</td>
<td>1.39±0.45</td>
<td>2.30±0.44</td>
<td>2.83±1.16</td>
</tr>
<tr>
<td>RS+PL(8)</td>
<td>3.72±0.91</td>
<td>0.64±0.35</td>
<td>1.47±0.48</td>
<td>2.20±0.46</td>
<td>2.77±1.29</td>
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<tr>
<td>RS+PL(9)</td>
<td>4.14±2.32</td>
<td>0.68±0.23</td>
<td>1.42±0.52</td>
<td>2.34±0.41</td>
<td>2.87±0.72</td>
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<tr>
<td>RS+PL(9-)</td>
<td>4.71±0.18</td>
<td>0.70±0.19</td>
<td>1.45±0.49</td>
<td>2.36±0.51</td>
<td>2.86±1.18</td>
</tr>
</tbody>
</table>

Figure 5.7: Confidence contours for Raymond-Smith + power law models for 600586p. The left hand panel shows power law photon index against Raymond-Smith temperature with $N_H$ fixed at the best fit value of $4.71 \times 10^{20}$ cm$^{-2}$. The centre panel shows power law photon index against $N_H$ with $kT$ fixed at the best fit value of 0.70keV. The right hand panel shows temperature against $N_H$ with the photon index fixed at the best fit value of 2.36. In both panels the contours mark 68%, 95% and 99% confidence intervals.
5.5 ROSAT HRI Pointed Observations

5.5.1 Source Detection

The ROSAT HRI is the only instrument to have observed NGC 4552 in repeated on-axis pointed observations. As such these observations should allow the best determination of source variability with the minimum of instrumental systematic problems affecting comparison.

<table>
<thead>
<tr>
<th>ID</th>
<th>Observation Date</th>
<th>SASS Version</th>
<th>Processing Date</th>
<th>OBIs</th>
<th>Exposure Livetime(s)</th>
</tr>
</thead>
<tbody>
<tr>
<td>600491h</td>
<td>9-11/07/1994</td>
<td>SASS7.3</td>
<td>30/09/1994</td>
<td>8</td>
<td>17996</td>
</tr>
<tr>
<td>600491h1</td>
<td>14-17/06/1995</td>
<td>SASS7.7</td>
<td>10/07/1995</td>
<td>13</td>
<td>29868</td>
</tr>
</tbody>
</table>

Table 5.7: ROSAT HRI observation summary.

The observation details listed in Table 5.7 show that both observations have SASS processing dates before March 1999 and, as described in Chapter 4 this includes an error in the calculation of the spacecraft aspect time leading to a broadening of the HRI+XRT point spread function. Before analysis these two HRI observations were corrected using the procedures developed by F. Primini (Mossman 1999) and the resulting event files were reduced using the standard EXSAS procedures. The X-ray contours from the centre of the 600491h image are plotted over the optical image in Fig 5.8 and the source detections in each observation are listed in Table 5.8. There is a ~30% drop in count rate for NGC 4552 between the two observations.
<table>
<thead>
<tr>
<th></th>
<th></th>
<th></th>
<th></th>
<th></th>
<th></th>
</tr>
</thead>
<tbody>
<tr>
<td><strong>600491h Detections</strong></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>4552</td>
<td>12°35′39.63″ +12°33′28.0″</td>
<td>796.5</td>
<td>23.43±1.17</td>
<td>0.32″</td>
<td>5.1±1.3</td>
</tr>
<tr>
<td>D</td>
<td>12°35′29.44″ +12°31′10.4″</td>
<td>13.3</td>
<td>1.09±0.30</td>
<td>3.69″</td>
<td>3.4±2.5</td>
</tr>
<tr>
<td>E</td>
<td>12°35′18.94″ +12°33′21.1″</td>
<td>12.0</td>
<td>0.91±0.28</td>
<td>5.35″</td>
<td>2.8±2.7</td>
</tr>
<tr>
<td>C</td>
<td>12°36′01.52″ +12°29′05.5″</td>
<td>39.1</td>
<td>1.80±0.36</td>
<td>6.77″</td>
<td>0</td>
</tr>
<tr>
<td>F</td>
<td>12°35′30.75″ +12°41′22.7″</td>
<td>10.9</td>
<td>0.96±0.30</td>
<td>8.16″</td>
<td>1.6±2.7</td>
</tr>
<tr>
<td>B</td>
<td>12°35′23.27″ +12°27′33.1″</td>
<td>11.2</td>
<td>1.56±0.41</td>
<td>9.04″</td>
<td>5.5±4.4</td>
</tr>
<tr>
<td><strong>600491h1 Detections</strong></td>
<td></td>
<td></td>
<td></td>
<td></td>
<td></td>
</tr>
<tr>
<td>4552</td>
<td>12°35′39.77″ +12°33′27.6″</td>
<td>570.1</td>
<td>16.67±0.78</td>
<td>0.29″</td>
<td>5.7±1.4</td>
</tr>
<tr>
<td>-</td>
<td>12°35′40.24″ +12°33′52.0″</td>
<td>14.4</td>
<td>2.34±0.35</td>
<td>0.30″</td>
<td>6.3±2.5</td>
</tr>
<tr>
<td>G</td>
<td>12°35′29.12″ +12°32′49.5″</td>
<td>10.9</td>
<td>0.48±0.16</td>
<td>2.96″</td>
<td>1.5±1.9</td>
</tr>
<tr>
<td>E</td>
<td>12°35′19.07″ +12°33′17.3″</td>
<td>80.4</td>
<td>1.96±0.28</td>
<td>5.32″</td>
<td>1.9±1.6</td>
</tr>
<tr>
<td>F</td>
<td>12°35′31.17″ +12°41′21.7″</td>
<td>12.3</td>
<td>0.81±0.23</td>
<td>8.11″</td>
<td>3.0±3.3</td>
</tr>
<tr>
<td>B</td>
<td>12°35′13.87″ +12°27′30.4″</td>
<td>14.6</td>
<td>0.95±0.24</td>
<td>8.96″</td>
<td>0</td>
</tr>
<tr>
<td>A</td>
<td>12°35′37.20″ +12°21′27.0″</td>
<td>25.6</td>
<td>2.02±0.38</td>
<td>12.18″</td>
<td>0</td>
</tr>
</tbody>
</table>

Table 5.8: ROSAT HRI source detection summary. None of the sources A-G have extension ML values above 2.3 and are therefore classified as point sources.
CHAPTER 5. THE SOFT X-RAY VARIABILITY OF NGC 4552

5.5.2 Extension

Radial profiles from the two HRI images were extracted around the detected source positions and fitted with a surface brightness model of a constant background component plus a $\beta$ model (e.g. Cavaliere & Fusco-Femiano 1976, Jones & Forman 1984) for the extended emission,

$$ S = S_{b_{\text{grad}}} + S_0 \left[ 1 + \left( \frac{r}{r_0} \right)^2 \right]^{-0.5-3\beta} $$

(5.1)

There is a very clear change in profile between the two observations with $r_0$ broadening, and $\beta$ steepening significantly in the second observation (Table 5.9). The model fit also worsens and is unacceptable ($\chi^2/\nu=15.85$). This
Table 5.9: Radial profile fit parameters to HRI data sets (See Fig 5.9). The degrees of freedom ($\nu$) were equal to 47 for each fit.

<table>
<thead>
<tr>
<th>Pointing</th>
<th>$S_{bgnd}$</th>
<th>$S_0$</th>
<th>$r_0''$</th>
<th>$\beta$</th>
<th>$\chi^2/\nu$</th>
</tr>
</thead>
<tbody>
<tr>
<td>600491h</td>
<td>81.8±2.5</td>
<td>11870±1526</td>
<td>4.01±0.47</td>
<td>0.52±0.02</td>
<td>1.09</td>
</tr>
<tr>
<td>600491h1</td>
<td>150.8±3.0</td>
<td>7689±3494</td>
<td>9.53±6.02</td>
<td>0.89±0.53</td>
<td>15.85</td>
</tr>
</tbody>
</table>

Figure 5.9: Beta models fitted to the 600491h (bottom) and 600491h1 (top) data. The 600491h data is shown with a beta model fit (dashed line). The 600491h1 data and fit (dotted line) have been shifted up by a factor 10 for clarity. Also overlaid in the 600491h1 data is a beta model (upper dot dashed line) using the 600491h values for $S_0$, $r_0$, and $\beta$ and the 600491h1 background. The best fit Gaussian ($\sigma$=3.3") to the HRI calibration PSF at 1keV is shown (lower dot dashed line) for reference.
suggests that the 600491h1 image suffers considerably more broadening from aspect error than the first observation.

Profiles were also extracted with centres on a grid of coordinates within the brightest pixel of each image but the best fit position was indistinguishable from that derived from the ML source detection shown in Table 5.8. This eliminates the possibility of broadening due to an off-centre fit.

Before examining the nature and cause of the count rate variability between the two observations it is important to determine the level of aspect error affecting the data.

5.5.3 Wobble correction.

The wobble detection procedures described in Chapter 4, for use on bright sources (Harris et al. 1998, Harris 1999) were applied to the data but no improvement was detected for either data set. This is not unexpected given that for sources of 0.023cts/sec and 0.016cts/sec the errors in source position when the data is split into 10 bins is of similar order to the expected wobble residuals.

The 'half phase' bin method was applied to both data sets therefore. The standard spacecraft wobble pattern (Fig 5.10) was carried out for both observations with the wobble for each OBI remaining in phase so no phase corrections were applied to the data. The roll stability and deviation from the wobble pattern are shown for each OBI in the two observations in Table
Figure 5.10: Both 491n00 (left) and 491a01 (right) showing a spacecraft wobble pattern with single phase throughout observations. This allows wobble correction to be applied to the whole events file without reregistering first. Only the declination component of the wobble is shown here as the RA component has only ~10% the declination amplitude. The parts of the attitude log plotted here that are not within the conventional wobble pattern are not included in the intervals used in the analysis.

5.10. Four OBIs in total had unacceptably high variation in roll and pointing during the observations. These occurred during 'settling' into the proper wobble pattern and all such time intervals were excluded from the subsequent analysis.

The half phase bin source detection routines were applied to both data sets. Both images show clear phase dependent variation (Fig 5.11) for the y coordinate with smaller amplitude variations in the x coordinate. The same data is shown as a connected track of points as the phase offset ($\phi$) is varied from 0 to 50% (Fig 5.12).
The amplitudes, \( A \), associated with the y coordinate are 3.2" and 3.75" for 600491h and 600491h1 respectively. These y coordinate corrections were applied to the photon events (the x coordinate was left un-altered) but no significant improvement over the uncorrected source extensions were achieved.
Figure 5.11: The detected coordinates (y, top; x bottom) in each of the two half phase binned images of 600491h (left) and 600491h1 (right) as the phase offset is varied. The error on each detected position is 0.8sky_pix.
Figure 5.12: The detected coordinates for 600491h (top) and 600491h1 (bottom) shown in Fig 5.11 plotted with successive offset positions connected. The error on each detected position is 0.8sky_pix in each axis.
### Table 5.10: Observation interval parameters for 600491h (top) and 600491hl (bottom). For Roll angle, RA and Declination, the mean of the standard deviations of the values in each of the 20 phase bins (as well as the value of the maximum s.d.) are listed. All units in sky_pix (0.5")

<table>
<thead>
<tr>
<th>OBI</th>
<th>Exposure (sec)</th>
<th>Roll RA</th>
<th>Roll Dec</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td></td>
<td>$\bar{\sigma}$ ($\sigma_{\text{max}}$)</td>
<td>$\bar{\sigma}$ ($\sigma_{\text{max}}$)</td>
</tr>
<tr>
<td>1</td>
<td>2608</td>
<td>14.8 (41.2)</td>
<td>13.0 (16.3)</td>
</tr>
<tr>
<td>2</td>
<td>7758</td>
<td>229.1 (234.8)</td>
<td>572.7 (646.3)</td>
</tr>
<tr>
<td>3</td>
<td>2236</td>
<td>25.3 (36.3)</td>
<td>10.6 (14.3)</td>
</tr>
<tr>
<td>4</td>
<td>2575</td>
<td>16.2 (37.0)</td>
<td>19.3 (77.4)</td>
</tr>
<tr>
<td>5</td>
<td>2642</td>
<td>16.5 (34.0)</td>
<td>17.6 (33.4)</td>
</tr>
<tr>
<td>6</td>
<td>1814</td>
<td>1.3 (40.3)</td>
<td>10.3 (25.4)</td>
</tr>
<tr>
<td>7</td>
<td>1668</td>
<td>25.6 (65.9)</td>
<td>14.5 (29.5)</td>
</tr>
<tr>
<td>8</td>
<td>2688</td>
<td>47.8 (105.4)</td>
<td>21.5 (50.6)</td>
</tr>
<tr>
<td>9</td>
<td>2770</td>
<td>10.9 (15.2)</td>
<td>32.7 (78.1)</td>
</tr>
<tr>
<td>10</td>
<td>1555</td>
<td>60.5 (103.8)</td>
<td>23.1 (39.5)</td>
</tr>
<tr>
<td>11</td>
<td>7530</td>
<td>71.6 (100.6)</td>
<td>955.8 (1036)</td>
</tr>
<tr>
<td>12</td>
<td>1985</td>
<td>76.0 (154.3)</td>
<td>30.7 (49.5)</td>
</tr>
<tr>
<td>13</td>
<td>1722</td>
<td>80.7 (111.2)</td>
<td>26.8 (48.2)</td>
</tr>
<tr>
<td>14</td>
<td>1987</td>
<td>78.7 (101.0)</td>
<td>29.1 (47.8)</td>
</tr>
<tr>
<td>15</td>
<td>7999</td>
<td>70.3 (93.7)</td>
<td>488.3 (1170)</td>
</tr>
<tr>
<td>16</td>
<td>738</td>
<td>68.3 (119.9)</td>
<td>15.7 (87.7)</td>
</tr>
<tr>
<td>17</td>
<td>862</td>
<td>61.0 (114.3)</td>
<td>23.8 (44.6)</td>
</tr>
<tr>
<td>18</td>
<td>1655</td>
<td>87.8 (179.2)</td>
<td>25.0 (57.3)</td>
</tr>
<tr>
<td>19</td>
<td>1213</td>
<td>47.4 (118.0)</td>
<td>15.5 (33.7)</td>
</tr>
<tr>
<td>20</td>
<td>13862</td>
<td>122.8 (142.4)</td>
<td>4385 (4956)</td>
</tr>
<tr>
<td>21</td>
<td>1108</td>
<td>70.9 (154.8)</td>
<td>17.6 (34.9)</td>
</tr>
</tbody>
</table>
5.5.4 OBI offset detection

To check whether the level of attitude solution errors during any of the Observation Intervals (OBIs) could have lead to the different radial profiles source detections were carried out on each OBI. For the two appropriate point source detections available within 5' offaxis (D and E), they were stacked with NGC 4552 such that their Gaussian centred peak coincided to attempt to improve the signal for source detection.

The positions for both NGC 4552 alone, and the galaxy stacked with source D and E (Table 5.8) are detected in each OBI (Table 5.11 and Fig 5.13). The significance of the OBI offset ($\sigma_{OBI}$) from the centroid is calculated based on the accuracy of the position (see caption for Table 5.11) from the source detection. To correct the image, OBIs with the largest $\sigma_{OBI}$ and detected counts above 50 were selected and the events during those periods were recentered at the centroid. The resulting detection parameters for these corrected images are shown in Table 5.12. No significant reduction in extension (i.e. EXT(ML) and FWHMs) is found from translating even just the most significant offset OBI back to the centroid. This implies that the source is too faint to detect a significant offset at this level and that much higher $\sigma_{OBI}$ would be required to be sure that an OBI is displaced.

The weighted mean OBI offset is calculated as $\frac{\sum_i (\Delta X_i) \cdot Counts_i}{\sum_i Counts_i}$ and the x and y components are 1.39 and 2.52 for the detections of NGC 4552 alone.
Although there are more integrated counts in the stacked detections and the weighted mean offset is increased, there are also no significant improvements in the detections of NGC 4552 in these images.

<table>
<thead>
<tr>
<th>#</th>
<th>Xcen</th>
<th>Ycen</th>
<th>ML</th>
<th>Counts</th>
<th>$\Delta X$</th>
<th>$\Delta Y$</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>35.22±2.14</td>
<td>15.74±2.36</td>
<td>154.7</td>
<td>0.36</td>
<td>0.68</td>
<td></td>
</tr>
<tr>
<td>2</td>
<td>35.27±1.15</td>
<td>17.80±2.39</td>
<td>154.4</td>
<td>0.41</td>
<td>1.66</td>
<td></td>
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<tr>
<td>3</td>
<td>35.30±1.14</td>
<td>17.65±2.32</td>
<td>10.6</td>
<td>0.87</td>
<td>1.27</td>
<td></td>
</tr>
<tr>
<td>4</td>
<td>35.72±1.14</td>
<td>17.99±2.32</td>
<td>59.6</td>
<td>1.38</td>
<td>4.37</td>
<td></td>
</tr>
<tr>
<td>5</td>
<td>36.79±1.13</td>
<td>18.04±2.32</td>
<td>70.0</td>
<td>3.09</td>
<td>4.19</td>
<td></td>
</tr>
<tr>
<td>6</td>
<td>34.37±1.12</td>
<td>17.42±2.29</td>
<td>30.3</td>
<td>0.70</td>
<td>0.16</td>
<td></td>
</tr>
<tr>
<td>7</td>
<td>36.28±3.34</td>
<td>19.33±3.36</td>
<td>36.9</td>
<td>5.03</td>
<td>3.28</td>
<td></td>
</tr>
<tr>
<td>8</td>
<td>33.47±1.14</td>
<td>14.58±1.18</td>
<td>146.0</td>
<td>1.45</td>
<td>2.98</td>
<td></td>
</tr>
<tr>
<td>9</td>
<td>31.90</td>
<td>19.90</td>
<td>3646.9</td>
<td>1382.9</td>
<td>2.02</td>
<td></td>
</tr>
</tbody>
</table>

Table 5 (c): Maximum likelihood detection parameters for each QBI using (top) NGC 4552 alone and (bottom) stacked with sources C and D line each QBI, the counts were all within a circle of 3.5° with 1σ error, the ML source detection parameters were searched within each QBI and errors from the source detection. Errors in $\Delta X, \Delta Y$ are the offsets from the QBI centroid position. The offsets for each QBI which were actually calculated by $\Delta X, \Delta Y$ are marked within (a) the centroid positions ($\bar{X}, \bar{Y}$) which were used to calculate (b) the weighted mean offset (see Equation 2). The weighted mean offset is a measure of the significance of the offset of each QBI from the overall centroid.
Table 5.11: Maximum likelihood detection parameters for each OBI using (top) NGC 4552 alone and (bottom) stacked with sources C and D. For each OBI, the centre Xcen,Ycen (units of 0.5") with 1-σ errors, the ML source detection and counts detected within each OBI are listed from the source detection. Column $\sigma_{OBI}$ is the offset from the OBI centroid $(\sqrt{\Delta X^2 + \Delta Y^2})/ (\sqrt{\sigma_{Xcen}^2 + \sigma_{Ycen}^2})$. The OBIs which were actually corrected by $\Delta X, \Delta Y$ are marked with †. (a) The centroid positions ($\frac{\sum_i (Xcen_i \cdot Counts_i)}{\sum_i Counts_i}$) (b) The weighted mean offset ($\frac{\sum_i (\Delta X_i \cdot Counts_i)}{\sum_i Counts_i}$). Nevertheless, $\sigma_{OBI}$ is a useful measure of the significance of the offset of each OBI from the overall centroid.
<table>
<thead>
<tr>
<th>Method</th>
<th>ML</th>
<th>cts</th>
<th>Xsky †</th>
<th>Ysky †</th>
<th>EXT(ML) †</th>
</tr>
</thead>
<tbody>
<tr>
<td>Uncorrected</td>
<td>769.5</td>
<td>420.67 ± 21.09</td>
<td>34.85 ± 0.6</td>
<td>15.97 ± 0.6</td>
<td>10.2 ± 2.6(302.2)</td>
</tr>
<tr>
<td>Primini Correction</td>
<td>795.4</td>
<td>420.04 ± 21.06</td>
<td>34.94 ± 0.6</td>
<td>16.39 ± 0.6</td>
<td>10.0 ± 2.6(286.8)</td>
</tr>
<tr>
<td>OBI-fix (4552 : OBI #1)</td>
<td>799.9</td>
<td>420.81 ± 21.08</td>
<td>34.94 ± 0.6</td>
<td>16.78 ± 0.6</td>
<td>10.0 ± 2.6(285.4)</td>
</tr>
<tr>
<td>OBI-fix (4552 : OBI #1,#8)</td>
<td>802.5</td>
<td>421.72 ± 21.11</td>
<td>34.97 ± 0.6</td>
<td>16.70 ± 0.6</td>
<td>10.0 ± 2.6(285.1)</td>
</tr>
<tr>
<td>OBI-fix (4552 : OBI #1,#2,#8)</td>
<td>805.3</td>
<td>419.89 ± 21.06</td>
<td>34.88 ± 0.6</td>
<td>16.40 ± 0.6</td>
<td>9.9 ± 2.6(280.3)</td>
</tr>
<tr>
<td>OBI-fix (4552+C+D : OBI #4)</td>
<td>797.9</td>
<td>419.03 ± 21.03</td>
<td>34.44 ± 0.6</td>
<td>16.62 ± 0.6</td>
<td>9.9 ± 2.6(286.2)</td>
</tr>
<tr>
<td>OBI-fix (4552+C+D : OBI #4,#8)</td>
<td>800.5</td>
<td>420.80 ± 21.08</td>
<td>34.51 ± 0.6</td>
<td>16.52 ± 0.6</td>
<td>10.0 ± 2.6(286.6)</td>
</tr>
<tr>
<td>OBI-fix (4552+C+D : OBI #1,4,#8)</td>
<td>806.2</td>
<td>420.59 ± 21.08</td>
<td>34.39 ± 0.6</td>
<td>16.85 ± 0.6</td>
<td>9.9 ± 2.6(283.2)</td>
</tr>
<tr>
<td>OBI-fix (4552+C+D : OBI #1,2,4,#8)</td>
<td>808.7</td>
<td>420.67 ± 21.08</td>
<td>34.14 ± 0.6</td>
<td>16.73 ± 0.6</td>
<td>9.9 ± 2.6(281.0)</td>
</tr>
</tbody>
</table>

<table>
<thead>
<tr>
<th>Method</th>
<th>Xcen†</th>
<th>Ycen†</th>
<th>Xsig†</th>
<th>Ysig†</th>
<th>XFWHM†</th>
<th>YFWHM†</th>
</tr>
</thead>
<tbody>
<tr>
<td>Uncorrected</td>
<td>34.73 ± 0.54</td>
<td>15.70 ± 1.04</td>
<td>10.17</td>
<td>11.78</td>
<td>23.95</td>
<td>27.73</td>
</tr>
<tr>
<td>Primini Correction</td>
<td>34.16 ± 0.53</td>
<td>15.86 ± 1.08</td>
<td>9.05</td>
<td>9.79</td>
<td>21.32</td>
<td>23.06</td>
</tr>
<tr>
<td>OBI-fix (4552 : OBI #1)</td>
<td>34.29 ± 0.62</td>
<td>16.42 ± 1.14</td>
<td>10.57</td>
<td>11.43</td>
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Table 5.12: OBI offset corrections in 600491h showing (top) ML Detection parameters and (bottom) Gaussian parameters. The OBIs chosen for correction are based on those in Table 5.11 with ≥ 50cts and σOBI > 1 (#2 was also accepted as a test). For the detections of the stacked events from NCG4552+C+D OBI #4 with 49.7cts is accepted also. † All coordinates are in units of 0.5". Xsky, Ysky and Xcen, Ycen are the detected offsets from the frame centre.
Figure 5.13: Positions of the detected OBI centres using (left) NGC 4552 detections alone and (right) NGC 4552 stacked with Srcs C and C. The area of each circle is proportional to the detected counts in that OBI. The OBI centroid is shown as a crossed square.
The procedures described for 600491h were repeated for 600491hl although only one source was available for stacking (E). The weighted mean OBI offsets of 3.66 and 4.84 sky_pix are larger than for 600491h and would appear to be a significant component of the image broadening.

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### CHAPTER 5. THE SOFT X-RAY VARIABILITY OF NGC 4552

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Table 5.13: Maximum likelihood detection parameters of M89 in each OBI using (top) M89 alone and (bottom) M89 stacked with source E (5.3 offaxis). OBI with All parameters are as listed in Table 5.11.
Figure 5.14: The original 600491h image (left) with the source+background regions superimposed on the central source (right) to allow analysis of the stacked sources. The Gaussian fitted centres of the two off-centre sources were translated to overlay on the Gaussian centre of NCG 4552.
Figure 5.15: 600491h1 Positions of the detected OBI centres using (left) M89 detections alone and (right) M89 stacked with Srcs A and B. The area of each circle is proportional to the detected counts in that OBI. The OBI centroid is shown as a crossed square. [XXX Image (a) produced with midwork/m89_OBI_srcs.prg (b) with midwork/stk_OBI_srcs.prg]
Table 5.14: Beta model fit parameters for OBIs #3, #7 and #12 in 600491h1 from Fig 5.12 and also after correcting for the detected wobble residual (lower three rows). For clarity the full image profile detection parameters are reproduced from Table 5.9. No significant improvement in each of the OBIs is seen after the wobble residual correction has been applied.

**Dewobbling individual OBIs in 600491h1**

The wobble correction determined in the previous section was used to attempt to correct the three longest OBIs. The radial profile fitting did not show any significant improvement on the total image.

**5.5.5 Variability**

The NGC 4552 source events for both data sets were binned over two wobble periods (800 seconds) into lightcurves and are both consistent with no change in count rate during the observations (see Fig 5.18).
Figure 5.16: Radial profiles of OBIs #3 (top), #7 (middle) and #12 (bottom) of 600491hl. The parameters of the beta model (green line) profile fits are shown in Table 5.14. Only OBI #12 shows any evidence of a central excess above the model. The background level (blue) was fitted to the profile between 100" and 200".
Figure 5.17: Contour plots of OBIs #3 (top left), #7 (top right) and #12 (bottom left) of 600491hl. Contours are at 7.2, 14.4, 28.8, 43.2, 57.6, 72 counts/arcmin²/ksec.
5.5.6 Spectral Analysis

As discussed in Chpt. 2.4.4 some limited spectral information is available from the HRI detector. I do not attempt to fit spectral models to the HRI data but instead search for indications of change in the pulse height amplitude (PHA) distribution which would indicate change in spectral components.

The events were extracted from within a 90″ × 90″ square box centred on the galaxy in each image. Backgrounds were extracted from a source free 90″ × 90″ box just to the North of the source box. The PHA distributions are plotted in Fig. 5.19. The use of the HRI for both spectral analysis and more importantly for X-ray flux determination, is complicated by the UV leak.
problem with the Lexan/Aluminium shield (Zombeck et al. 1997, Berghöfer, Schmitt & Hünsch 1999, Barbera et al. 2000). However, the UV leak and particle backgrounds are expected to contribute principally to channels #1-3, and in this case the total contributions for these channels are 24% and 21% of the total counts for 600491h and 600491h1 respectively. So it is most likely that the contribution of UV emission from NGC 4552 to the total count rate will be below that level.

The PHA distributions in Fig 5.19 show the usual concentration of HRI counts in channels 5-8. The S-profile in the differences between the normalised distributions of the two sources clearly indicates the shifting to lower energies of the 600491h1 data. The ongoing reduction in the gain of the HRI shifts the mean PHA channel downwards at a rate of $\sim 0.5$ channel/year (Prestwich et al. 1996). Therefore, to check if this shift is sufficient to explain the change in PHA distribution, (and since the two observations were taken 11.2 months apart), the differences are also calculated for the data with a 0.5 channel shift (see row 491h1(c) in Table 5.15).

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Table 5.15: Spectral properties for the ROSAT HRI detections. The last rows show the source and background parameters when the data have been corrected by 0.5 channels.
Figure 5.19: The binned raw PHA distributions for events extracted within 90" x 90" square boxes centred on the NGC 4552 detected position for 600491h (top left) and 600491h1 (top right). The backgrounds were extracted from a source free 90" x 90" box just to the North of the source box. The lower pair of graphs show the difference (491h minus 491h1) between the background subtracted normalised PHA distributions. The differences (bottom left) show the clear 'S' profile due to the shift in PHA distribution. When the nominal 0.5 PHA channel shift has been added to the 600491h1 data (bottom right), the only clear feature remaining is the excess in channel #1 due to the varying background.
As can be seen from the bottom right graph in Fig 5.19 and from the similarity in fractional hard and soft source counts in Table 5.15, the differences in the two distributions can be principally attributed to the changing HRI gain.
5.6 Einstein

In July 1979 Einstein observed NGC 4552 at an off-axis angle of 9.7' for 8426 seconds with the IPC (0.2-3.5 keV). The corresponding image is shown in Fig. 5.20. The galaxy was detected at a count rate of 35.16±2.7 cts/ksec (Fabbiano et al. 1992). Previous analyses of this data (Forman, Jones & Tucker 1985, Kim, Fabbiano & Trinchieri 1992) presented various spectral models for NGC 4552.

Figure 5.20: Full energy band (0.3-3.5keV) Einstein IPC image (left) showing NGC 4552 within the 1° square inner FOV. (right) The IPC intensity contours for NGC 4552 overlaid on the DSS optical image. The contours show the symmetrical point source at the centre and the North-South extension at larger radii.

Forman, Jones & Tucker (1985) determined the M89 flux using a fixed 1keV thermal model. The best fit model from Kim, Fabbiano & Trinchieri (1992) for an absorbed Raymond Smith (RS) thermal plasma gives kT=1.23keV
and $N_H = 1.7 \times 10^{20} \text{cm}^{-2}$ for $\chi^2/\nu = 21.1/8$ although an absorbed power law (PL) of $\Gamma = 2.7$ and $N_H = 5 \times 10^{21} \text{cm}^{-2}$ fits almost as well. Due to the limited number of counts (280), more complicated models are not well constrained.

Figure 5.21: Einstein IPC spectrum showing the fitted model (solid line) of a $0.70\text{keV}$ Raymond Smith thermal plasma with an absorbing column of $N_H = 4.71 \times 10^{20} \text{cm}^{-2}$.

Therefore for flux calculation, I have adopted the best fit RS+PL model from the 1993 PSPC data and fit it to the Einstein data. I extracted the source events from a box $5'$ square centred on the source and a background from an annulus of $10' \times 10'$ around the source. During fitting only the PL and RS component normalisations were allowed to vary. For model (a) in Table 5.19, the relative normalisations of the PL and RS components were fixed to the 1993 PSPC model and only the total model normalisation could vary. For model (b) the normalisations of the power law and Raymond-Smith components were allowed to vary independently. Here, the power law normalisation goes to zero and $\chi^2/\nu$ improves from 1.47 to 1.07. This fit to the data is shown in Fig 5.21.
CHAPTER 5. THE SOFT X-RAY VARIABILITY OF NGC 4552

5.7 EXOSAT

The EXOSAT satellite observed the Virgo cluster in a set of pointed observations during July 1983. The data from the non-imaging Medium Energy (2-30keV) instrument are dominated by emission from M87 (71' away from NGC 4552) and the Virgo cluster gas and are not used here. The two Low Energy (LE) telescopes (0.05-2.5 keV) with their associated Channel Multiplier Array (CMA) detectors imaged the NGC 4552 region with various filters in thirteen observations during this period. The CMA detector was read out using a resistive plate charge collector with four readout electrodes which give rise to the stuctures visible in Fig 5.22.

I carried out standard source detection on the binned events images with the XIMAGE package. Source detections and 3σ upper limits are given in Table 1. All observations of NGC 4552 apart from a03574 were too far off-axis to yield reliable results. No upper limits are given for seven observations with off-axis angles greater than 57' due to detection artifacts near the edge of the field. The a03574 LE2 observation (15.3' offaxis) with the 3000Å Lexan filter (see Fig 3.3) provides the most stringent upper limit on the NGC 4552 flux for this epoch. The detection in a03604 is associated with artifacts in the background due to the geometry of the readout electronics which render this value unreliable. The non-detection of NGC 4552 at the upper limit of 2cts/ksec is consistent with the observation of NGC 1399 which was detected at 5cts/ksec, 5' off-axis, in an exposure time of 24ksec using the CMA/LE combination and the same filter (Mason & Rosen 1985).
Figure 5.22: EXOSAT LE/CMA image a03604 showing the artifacts in the background near the NGC 4552 position (white circle bottom left) which lead to a false detection. The readout electrode positions are at top, bottom, left and right.

<table>
<thead>
<tr>
<th>Date</th>
<th>Frame ID</th>
<th>Detector</th>
<th>Time (ksec)</th>
<th>Count Rate (/ksec)</th>
<th>Off-axis angle</th>
</tr>
</thead>
<tbody>
<tr>
<td>14/7/1983</td>
<td>a03552</td>
<td>LE2/3Lx</td>
<td>7.3</td>
<td>&lt;19.1</td>
<td>43.6'</td>
</tr>
<tr>
<td>15/7/1983</td>
<td>a03558</td>
<td>LE1/3Lx</td>
<td>4.3</td>
<td>&lt;28.0</td>
<td>42.9'</td>
</tr>
<tr>
<td>15/7/1983</td>
<td>a03559</td>
<td>LE2/Al/P</td>
<td>3.5</td>
<td>&lt;33.2</td>
<td>43.7'</td>
</tr>
<tr>
<td>15/7/1983</td>
<td>a03574</td>
<td>LE2/3Lx</td>
<td>27.8</td>
<td>&lt;2.0</td>
<td>15.3'</td>
</tr>
<tr>
<td>15/7/1983</td>
<td>a03604</td>
<td>LE2/3Lx</td>
<td>48.8</td>
<td>(13 ± 2)</td>
<td>49.0'</td>
</tr>
<tr>
<td>15/7/1983</td>
<td>a03605</td>
<td>LE1/3Lx</td>
<td>16.3</td>
<td>&lt;13.8</td>
<td>51.3'</td>
</tr>
</tbody>
</table>

Table 5.16: EXOSAT LE observations of NGC 4552 with off-axis angles <57'. The count rate refers to the 3σ upper limit (or the actual detected count rate for a03604) in each image. The telescopes LE1 and LE2 are both used with the CMA detector and either the 3000Å Lexan (3Lx) or Aluminium/Parylene (Al/P) filters.
I converted the EXOSAT upper limit count rate of 2 cts/ksec to a flux in the 0.05-2.5 keV LE/CMA band using the FTOOL command 'exopha' and calculated the flux using 'model/flux' command within XSPEC with the best fit model parameters as determined for the 600586p ROSAT PSPC observation in Table 5.6. This gives a flux upper limit of \(7 \times 10^{-12} \text{erg.s}^{-1} \text{.cm}^{-2}\) for the same model in the ROSAT (0.1-2.4 keV) band.

Table 5.6: Spectral model components from Valtti & Mushotzky (1993)

<table>
<thead>
<tr>
<th>Model</th>
<th>$kT$ (keV)</th>
<th>$N_{H}$ (cm$^{-2}$)</th>
<th>$A_{p}$</th>
<th>$A_{t}$</th>
<th>$N_{e}$</th>
<th>$N_{H}$</th>
<th>$A_{bol}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>FG</td>
<td>0.22</td>
<td>3.6</td>
<td>1.2</td>
<td>3.6</td>
<td>2.5</td>
<td>1.2</td>
<td>3.6</td>
</tr>
<tr>
<td>RS</td>
<td>0.35</td>
<td>1.5</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
</tr>
<tr>
<td>RS+BD</td>
<td>0.35</td>
<td>1.5</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
</tr>
<tr>
<td>DBBD+FL</td>
<td>0.35</td>
<td>1.5</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
<td>0.1</td>
</tr>
</tbody>
</table>

CM92 fitted a radial profile from the ROSAT image with a \(\beta\)-model plus a Gaussian to model the central emission. They found that the Gaussian component of the profile contributes \(1\%\) of the counts and therefore they used that fraction of the total flux as their estimate for the emission from the nucleus. Due to the high flux offset of CM92, they underestimated it by that amount.
5.8 ASCA

ASCA observed M89 in June 26th, 1995 (just 12 days after the 600491hl ROSAT HRI observation) with the SIS and GIS instruments for 30.5 and 32.0ksec respectively. For comparison with the ROSAT data the spectral analysis of the ASCA data by Colbert & Mushotsky (1999) (CM99) is used (Table 5.17).

<table>
<thead>
<tr>
<th>Model</th>
<th>$kT^a$ (keV)</th>
<th>$\Lambda^b_{RS}$</th>
<th>$\Gamma^c$</th>
<th>$\Lambda^d_{PL}$</th>
<th>$N_H^e$</th>
<th>Abund.$^f$</th>
<th>$\chi^2_\nu$</th>
</tr>
</thead>
<tbody>
<tr>
<td>PL</td>
<td>-</td>
<td>-</td>
<td>2.24$^{+0.44}_{-0.26}$</td>
<td>2.70</td>
<td>0.558$^{+0.07}_{-0.088}$</td>
<td>404</td>
<td>248</td>
</tr>
<tr>
<td>RS</td>
<td>0.76$^{+0.81}_{-0.71}$</td>
<td>21.1</td>
<td>-</td>
<td>-</td>
<td>2.23$^{+3.35}_{-1.45}$</td>
<td>0.10$^{+0.16}_{-0.07}$</td>
<td>454</td>
</tr>
<tr>
<td>RS+RL</td>
<td>0.69$^{+0.74}_{-0.64}$</td>
<td>1.16</td>
<td>2.03$^{+2.39}_{-1.64}$</td>
<td>2.33</td>
<td>7.71$^{+14.43}_{-2.18}$</td>
<td>1.17$^{+5.00}_{-0.57}$</td>
<td>243</td>
</tr>
<tr>
<td>DBB+PL</td>
<td>0.065$^{+0.064}_{-0.067}$</td>
<td>3.2e12</td>
<td>2.70$^{+2.49}_{-2.93}$</td>
<td>6.14</td>
<td>14.8$^{+14.1}_{-15.6}$</td>
<td>263</td>
<td>246</td>
</tr>
</tbody>
</table>

Table 5.17: Spectral model components from Colbert & Mushotsky (1999). The relatively high value of $N_H$ needs to be examined carefully since it is strongly dependent on the lower energy channels of the ASCA detector which are subject to some response uncertainty. Notes: (a) Temperature for the Raymond-Smith plasma model or surface temperature of the inner radius of the accretion disk ($kT_m$), (b) The RS normalisation is in units of (c) powerlaw slope, (d) The PL normalisation is in units of $10^{-23}$ photons cm$^{-2}$s$^{-1}$keV$^{-1}$, (e) Absorption column in units of $10^{21}$cm$^{-2}$ (f) Elemental abundances scaled to solar.

CM99 fitted a radial profile from the 600491hl image with a $\beta$-model plus a gaussian to model the nuclear emission. They found that the Gaussian component of the profile contributed 17% of the counts and therefore they used that fraction of the total flux as their estimate for the emission from the nucleus. Due to the significant effects of OBI and wobble residuals on that
CHAPTER 5. THE SOFT X-RAY VARIABILITY OF NGC 4552

observation, it is unsafe to trust radial profile fitting of the uncorrected data. Instead of the 17\% flux ($1.232 \times 10^{-13}$ erg s$^{-1}$ cm$^{-2}$) quoted in CM99, the full 100\% of the 0.1-2.4keV band flux for their best RS+PL model fit (see Table 5.19) is used yielding $F_X = 7.7 \times 10^{-13}$ erg s$^{-1}$ cm$^{-2}$.

5.9 Chandra

A 54ks Chandra/ACIS-S observation (OBS ID 2072) of NGC 4552 was made on April 22nd, 2001 and became public in April 2002. A full analysis is beyond the scope of the current work but I present a preliminary analysis to determine its consistency with the earlier X-ray data. I have used the source detections from the XASSIST (www.xassist.org) processing of the Chandra data set. The ACIS image (Fig 5.23) clearly shows the North-South elongated extended hot ISM emission detected with ROSAT HRI.

The resolving power of Chandra allows many individual point sources to be detected which have not been detected previously (Fig 5.24). However, the densest region around the nucleus including the positions of the HST UV flares within a few 0.1\" of the nucleus remains unresolved. The distribution of luminosities (assuming $\Gamma=1.8$, $N_H$ galactic) of all point sources detected with the XASSIST pipeline is shown in Fig 5.25. These are consistent with the point source population in the field being low mass X-ray binaries (LMXBs) (Irwin, Sarazin & Bregman 2002) although further from the halo of the galaxy it is likely that there are background AGNs also detected.
Figure 5.23: The DSS optical image of NGC 4552 (top) with the X-ray contours from the ACIS S3 CCD and the binned X-ray image (bottom) of the central square arcminute around the galaxy.
As an added check of whether systematic effects are an explanation of the variation between the ROSAT HRI observations a correlation of all sources detected with the ROSAT HRI above ML=10 was made with the detections for the 1993 PSPC and the Chandra observations (Table 5.19). Positional matches were accepted within 5" of a Chandra source for HRI and 30" for PSPC. The automatic spectral fitting by XASSIST of an absorbed powerlaw give $\Gamma$ in the range 1.6-2.0 with little absorption. Since the absorption is poorly constrained in these fits and is important for conversion to expected ROSAT HRI count rates, these values must be treated with some caution. The sources D and E have been identified as IXO candidates by Colbert & Ptak (2002) (their numbers #64 and #63 respectively), since they have 2-10keV luminosities above $10^{39}$erg.s$^{-1}$ (at the distance of NGC 4552) and lie
within two optical $R_{25}$ radii of the centre of the galaxy.

Of the seven HRI objects, four were significantly fainter in the second observation with two significantly brighter and one constant. This is consistent with the reduction in count rate for NGC 4552 over one year between the two HRI observations may be due to systematic problems related to the degraded aspect solution though the result is not statistically significant in itself.

The hot ISM component has major and minor diameters of $30^\prime\prime$ and $24^\prime\prime$ (2.4kpc and 1.6kpc) respectively (Fig 5.23). This is somewhat larger than the warm H$_\alpha$ component which also shows some asymmetry (Macchetto et al. 1996). The X-ray emission is similar in extent to the ISM components detected in two other giant ellipticals NGC 1553 and NGC 5846 (Trinchieri & Goudfrooij 2002).

![Figure 5.25](image)

Figure 5.25: Distribution of luminosities in the point sources detected in the Chandra/ACIS image. The single source shown in the rightmost bin of the histogram is actually the central source at $6\times10^{39}\text{erg.s}^{-1}$. 
<table>
<thead>
<tr>
<th>HRI Id</th>
<th>Position (2000)</th>
<th>600586p (^1)</th>
<th>600491h</th>
<th>600491h1</th>
<th>Chandra (^1) (2-10keV)</th>
</tr>
</thead>
<tbody>
<tr>
<td>Src A</td>
<td>12h35m36.6s,12d21m27.0s</td>
<td>2.47±0.58 (0.80)</td>
<td>≤1.41</td>
<td>2.09±0.38</td>
<td>-</td>
</tr>
<tr>
<td>Src B</td>
<td>12h35m13.9s,12d27m30.5s</td>
<td>4.73±0.67 (1.53)</td>
<td>1.48±0.41</td>
<td>0.91±0.24</td>
<td>2.28±0.22 (1.51)</td>
</tr>
<tr>
<td>Src C</td>
<td>12h36m01.5s,12d29m06.2s</td>
<td>2.84±0.51 (0.92)</td>
<td>1.78±0.36</td>
<td>≤0.70</td>
<td>-</td>
</tr>
<tr>
<td>Src D</td>
<td>12h35m29.5s,12d31m07.8s</td>
<td>5.54±0.82 (1.80)</td>
<td>1.08±0.30</td>
<td>≤0.41</td>
<td>3.33±0.26 (1.78)</td>
</tr>
<tr>
<td>Src E</td>
<td>12h35m19.1s,12d33m17.4s</td>
<td>2.19±0.52 (0.71)</td>
<td>0.90±0.28</td>
<td>1.96±0.28</td>
<td>4.96±0.31 (2.65)</td>
</tr>
<tr>
<td>Src F</td>
<td>12h35m31.2s,12d41m21.8s</td>
<td>≤2.38 (&lt;0.77)</td>
<td>0.95±0.30</td>
<td>0.81±0.23</td>
<td>3.61±0.29 (1.92)</td>
</tr>
<tr>
<td>Src G</td>
<td>12h35m29.1s,12d32m49.5s</td>
<td>≤0.34 (&lt;0.11)</td>
<td>0.48±0.16</td>
<td>≤0.32</td>
<td>-</td>
</tr>
</tbody>
</table>

Table 5.18: Count rates for all the ROSAT HRI source detections (above ML=10) with the 1993 ROSAT PSPC observation and the 2000 Chandra ACIS observation. 95.4% confidence upper limits are given for PSPC and HRI non-detections. Although source G is not shown in Table 5.8, it was detected ML=9.9 I have included it here as a detection rather than give an upper limit. Sources E and F are identified with B\(_\text{mag}\) 20.0 and 16.2 stars from USNO A2.0. The Chandra non-detections are due to the sources being not in the field of view or (for Src G) in the gap between chips. \([1]\) The equivalent HRI count rate (0.1-2.4keV) is given in parentheses for an absorbed powerlaw (\(\Gamma=1.8, N_H=2.57\times10^{20}\text{cm}^{-2}\)). For a thermal bremsstrahlung spectrum (\(kT=5\text{keV}, N_H=2.57\times10^{20}\text{cm}^{-2}\)) count rates are \(~4\%\) higher for PSPC and \(~20\%\) lower for Chandra.
5.10 Discussion

5.10.1 Long-term X-ray Flux Variability

The integrated soft X-ray flux from NGC 4552 has remained within a factor of two of $2 \times 10^{-12} \text{erg.s}^{-1}.\text{cm}^{-2}(0.1-2.4\text{keV})$ with variations seen regularly over the last 20 years of observation. Although there is no clear indication of systematic changes between different instruments/satellites, some of this measured variation is likely to be due to calibration differences between observations.

The HRI observations show a $\sim 30\%$ drop in the count rate of NGC 4552. The difference in PSF between the two observations is most likely due to the greater OBI offsets in the 600491h1 observation with some contribution from a greater wobble residual for that data also. There is no indication for count rate variation during either observation or for systematic variation of other sources detected in the field to explain the $\sim 30\%$ drop in count rate in the year between the observations.

From the Chandra data it appears that $\sim 25\%$ of the soft X-ray emission may be due to the central source with the remainder due to the thermal ISM component and the LMXB population. It would require a large fraction of the LMXB population changing from high state to low state together to account for the $30\%$ drop in HRI count rate over one year. This is unlikely and further strengthens the case for instrumental effects being responsible for the HRI variation.
<table>
<thead>
<tr>
<th>Date</th>
<th>Instrument/Detector</th>
<th>Γ</th>
<th>kT (keV)</th>
<th>N&lt;sub&gt;H&lt;/sub&gt; (10&lt;sup&gt;20&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;)</th>
<th>RS/PL</th>
<th>f&lt;sub&gt;X&lt;/sub&gt;(0.1-2.4 keV) 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</th>
<th>L&lt;sub&gt;X&lt;/sub&gt;(0.1-2.4 keV) 10&lt;sup&gt;40&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt;</th>
</tr>
</thead>
<tbody>
<tr>
<td>1979-07-01</td>
<td>Einstein/IPC</td>
<td>[a]2.36</td>
<td>0.70</td>
<td>4.71</td>
<td>0.49</td>
<td>1.99 ± 0.07</td>
<td>6.60±0.24</td>
</tr>
<tr>
<td></td>
<td></td>
<td>[b]2.36</td>
<td>0.70</td>
<td>4.71</td>
<td>∞</td>
<td>1.01 ± 0.06</td>
<td>3.40±0.20</td>
</tr>
<tr>
<td>1983-07-15</td>
<td>EXOSAT/LE</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>1.01 ± 0.06 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</td>
<td>&lt;23.64</td>
</tr>
<tr>
<td>1990-07</td>
<td>ROSAT/PSPC(ASS)</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>2.52 ± 0.31 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</td>
<td>8.51±1.05</td>
</tr>
<tr>
<td>1991-12-15</td>
<td>ROSAT/PSPC</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>1.44 ± 0.09 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</td>
<td>4.86±0.30</td>
</tr>
<tr>
<td>1992-07-06</td>
<td>ROSAT/PSPC</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>3.43 ± 0.31 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</td>
<td>11.58±1.04</td>
</tr>
<tr>
<td>1992-12-19</td>
<td>ROSAT/PSPC</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>1.81 ± 0.04 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</td>
<td>6.11±0.14</td>
</tr>
<tr>
<td>1993-06-30</td>
<td>ROSAT/PSPC</td>
<td>2.36</td>
<td>0.70</td>
<td>4.71</td>
<td>0.49</td>
<td>3.06 ± 0.03 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</td>
<td>10.33±0.10</td>
</tr>
<tr>
<td>1994-07-09</td>
<td>ROSAT/HRI</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>1.95 ± 0.04 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</td>
<td>6.59±0.14</td>
</tr>
<tr>
<td>1995-06-14</td>
<td>ROSAT/HRI</td>
<td></td>
<td></td>
<td></td>
<td></td>
<td>1.39 ± 0.03 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</td>
<td>4.69±0.10</td>
</tr>
<tr>
<td>1995-06-26</td>
<td>ASCA/GIS+SIS</td>
<td>2.03</td>
<td>0.69</td>
<td>77.1</td>
<td>0.52</td>
<td>0.77 ± 0.05 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</td>
<td>2.60±0.17</td>
</tr>
<tr>
<td>2000-04-22</td>
<td>Chandra/ACIS-S</td>
<td>2.0</td>
<td></td>
<td>2.57</td>
<td>-</td>
<td>0.50 ± 0.01 10&lt;sup&gt;-12&lt;/sup&gt; erg s&lt;sup&gt;-1&lt;/sup&gt; cm&lt;sup&gt;-2&lt;/sup&gt;</td>
<td>1.69±0.03</td>
</tr>
</tbody>
</table>

Table 5.19: Summary of all X-ray detections of NGC 4552. Observations for which no spectral fitting was done have fluxes calculated using the 1993 PSPC best fit model. RS/PL gives the relative normalisation of the Raymond-Smith and powerlaw components. For the Einstein spectral fits, [a] the 1993 PSPC model was used and in [b] the RS/PL normalisation was allowed to vary. The ASCA RS+PL model is from Table 5.18 using the flux from Colbert & Mushotsky (1999).
5.10.2 Spectral variability

There is no unambiguous evidence in favour of spectral variation on NGC 4552. The interpretations of small scale changes in relative band fluxes and spectral model components are complicated by systematic effects due to the variety of instruments, off-axis angles, and observing modes employed. In particular the off-axis ROSAT PSPC detections listed in Table 5.3 are strongly affected by changes in background (due to changes in the local particle flux) and off-axis angle.

Despite its relative insensitivity to spectral information the analysis of the pulse height distributions does not show any significant spectral change between the two ROSAT HRI observations. If there was a significant hardening from the first to the second observation then it could have been argued that the excess brightness in the first observation was due to a transient soft component. This is not strongly indicated by the HRI hardness ratio data.

5.10.3 Source Extention

All four observatories that detected NGC 4552 show that some level of extension is present. The strongest constraints placed on the extension are from the ROSAT HRI and Chandra ACIS. The Einstein, ROSAT PSPC and HRI, ASCA and most recently Chandra observations all show significant extension in the X-ray distribution. This extension is unresolved but clearly detected in the Einstein and ROSAT observations.
The change in the detected source extension in the ROSAT HRI images is largely explained by the aspect solution problems though no useful limit can be placed on the possible variation of LMXBs close to the centre of the nuclear peak contributing to this change.

The good fit of a beta model to the 600491h aspect time corrected data suggests that this profile is close to the true underlying distribution and certainly represents the best limit on the X-ray nuclear profile during the UV flare. A detailed comparison of this observation with the Chandra light distribution is also needed to complete the picture of variability in the nuclear region and its surroundings.

5.10.4 Was the UV flare caused by stellar disruption?

It seems clear from the broad lines measured in the FOC spectrum of the flare that there is an accretion structure orbiting the central black hole in NGC 4552. Whether this accretion was fuelled by stellar disruption or another mass infall source is not resolved. The strongest signature of a stellar disruption event is the decay curve being consistent with the predicted mass fallback rate. If the decay in the UV output between 1993 and 1996 was to be explained purely by a reduction in mass fall back rate, the stellar disruption event would have had to occur prior to 1991 leading to the question of why a brighter source was not observed in the nucleus in the July 1991 HST/FOC observation.

Using the mass estimate for NGC 4552 of $M = 4.66 \times 10^8 M_\odot$ (Magorrian et
Figure 5.26: The UV FOC data from Cappellari et al. (1999) plotted at the effective wavelength for each filter. The black body spectra shown at $kT=0.03\text{keV}$, $0.05\text{keV}$ and $0.07\text{keV}$ (respectively $T=3.5$, 5.8 and $8.1\times10^{6}\text{K}$) are normalised to fit the June 1993 PSPC observation at 23cts/ksec in the $0.1-0.5\text{keV}$ band. The UV data have errors of $\sim10\%$.  

al. 1998) gives $L_{Edd} \sim 6\times10^{46}\text{erg.s}^{-1}$. The exact choice of spectral model for the flare component does not significantly affect the conclusion that the flare is significantly sub-Eddington. It must be borne in mind also that these luminosities are upper limits given that much of the soft band emission in the 600586p observation can be explained by the power law and thermal components already fitted.

Even though the UV fluxes associated with the black body spectra are close to
### Table 5.20: Predicted flare luminosities from 600586p soft band counts for the three models shown in Fig 5.26. These values assume that the soft band count rate in the 1993 PSPC observation (0.023cts/sec) was due to a black body component associated with the flare.

<table>
<thead>
<tr>
<th>kT(_{bb}) (keV)</th>
<th>(f_X(0.1-0.5)) (10^{-13}\text{erg.s}^{-1}\text{.cm}^{-2})</th>
<th>(L_X) (10^{40}\text{erg.s}^{-1})</th>
<th>(L_{\text{bol}}/L_X) (10^{-7})</th>
<th>(L_{\text{bol}}/L_{\text{Edd}}) (0.1-0.5)</th>
<th>(10^{-7})</th>
</tr>
</thead>
<tbody>
<tr>
<td>0.03</td>
<td>9.24</td>
<td>3.12</td>
<td>1.87</td>
<td>9.7</td>
<td></td>
</tr>
<tr>
<td>0.05</td>
<td>4.73</td>
<td>1.60</td>
<td>1.24</td>
<td>3.3</td>
<td></td>
</tr>
<tr>
<td>0.07</td>
<td>4.50</td>
<td>1.52</td>
<td>1.18</td>
<td>3.0</td>
<td></td>
</tr>
</tbody>
</table>

Given the low limits on the luminosity of the flare of \(\sim 10^{40}\text{erg.s}^{-1}\) (or considerably lower if the flare is actually limited to the UV), the mass budget required to explain the luminosity by accretion is given by \(\dot{M} = L/\epsilon c^2\), or,

\[
\dot{M} = 1.6 \times 10^{-6} \left( \frac{L}{10^{40}\text{erg.s}} \right) \left( \frac{0.1}{\epsilon} \right) M_\odot\text{.yr}^{-1}
\]  

(5.2)
Over the lifetime of the flare from 1991-1996 the limit on the total mass infall would be $\sim 10^{-5} M_\odot$ assuming $\epsilon=0.1$. Such low masses may be supplied as suggested from the stripping of the atmosphere of a giant star (Renzini et al. 1995, Syer & Ulmer 1999) during a pericentre passage a few times $r_s$ from the BH.

A further source of mass from stellar disruption around very massive ($M_{BH}>10^8 M_\odot$) black holes which has not previously been discussed is mass loss during tidal capture of low mass stars.

The density of non-giant stars means that the disruption radius around a massive BH will be within the event horizon (see Chapter 2.2). Stars passing outside the disruption radius, $r_T$, may lose enough orbital angular momentum through tidal dissipation to become bound to the BH and orbital decay over subsequent pericentre passages will ultimately lead to the star passing within $r_T$ (Novikov, Pethick & Polnarev 1992). Even this process may not lead to disruption before advection around a BH as massive as appears to be in NGC 4552.

Stars passing far enough out from the BH not to be tidally captured and advected may nevertheless have tidal energy input to cause mass loss. The short time ($\sim 1hr$, approximately independent of BH mass) during which the changes in position angle and magnitude of the tides occur mean that significant energy is input on a timescale short compared to the star’s ability to radiate it at $L_{Edd}$. If the energy input is small the star may escape intact with a temporary radial expansion ‘pulse’ as the extra energy is radiated.
away. For higher energy input (in particular if shocks form inside the star), material in the outer layers of the star may be ejected in a hot envelope as the star expands (or shocks move up through the star to the surface). If such a shell was ejected with any significant energy, then some of the mass would become bound to the BH even if the centre of mass of the star remained unbound.

This process has been considered (Carter & Luminet 1982, Carter & Luminet 1983) for the case of a solar mass star passing within $r_T/3$ where they propose that the shocks forming would trigger much higher nucleosynthesis rates and possibly a helium flash which would effectively unbind the star. Even for the weaker tidal interaction outside the tidal capture radius ($\sim 3r_T$) mass loss of the order $10^{-5}$-$10^{-4}$ $M_\odot$ may be possible from solar density stars in particular if they are also spun up during approach to pericentre.

Alternatively, the mass infall source may not be related to stellar disruption. A low level infall of material from the surrounding molecular clouds or other gas from the ISM (Ciotti et al. 1991) could provide mass inflow although it would be retarded by heating if significant luminosity is generated in the accretion structures (Ciotti & Ostriker 2001, Bringenti & Mathews 2002). Any such scenario would not have the decay law constraints of the stellar disruption mass fallback since the mass would not necessarily be distributed as equal mass fractions per interval of orbital energy.

The other clear central spike candidates observed (Cappellari et al. 1999), NGC 2681 and NGC 1399 have not so far appeared to have the same class of
flares as NGC 4552. The steep central brightness profile of NGC 2681 peaking at \( \sim 14\text{mag/arcsec}^2 \) compared to \( \sim 15.5\text{mag/arcsec}^2 \) for the flare in NGC 4552 means that even if there was such a flare at that brightness in NGC 2681 then it would be almost impossible to detect (Cappellari et al. 1999, Cappellari et al. 2001).

Cappellari et al. (1999) have searched the FOC archive for other UV peaks in the nuclei if galaxies. They find that up to 25\% of early type galaxies may have central excesses above a 'Nuker'\(^2\) law (Lauer et al. 1995) but to date, NGC 4552 is the only nucleus showing any variability.

One interesting object is NGC 4374 (M84, 3C271.1) which as well as having a similar parsec-scale radio jet to NGC 4552 (Nagar et al. 2002) (see below), shows a clear nuclear peak in the single HST/FOC (F370LP filter) observation made in March 1991. Unfortunately there are no follow-up FOC observations to determine whether the nuclear emission at UV wavelengths shows similar variations to NGC 4552.

### 5.10.5 Alternatives to transient accretion

The two most compelling factors in favour of transient accretion for the UV event are the broad line emission and the very close alignment of the flare

\(^2\)The 'Nuker' profile (a label adopted by the literature after papers from Lauer's research group) is used to fit the radial profiles of elliptical galaxies and takes the form:

\[
I(r) = I_b \, 2^{\frac{\beta-\gamma}{\alpha}} \left( \frac{r}{r_h} \right)^{-\gamma} \left[ 1 + \left( \frac{r}{r_h} \right) ^{\alpha} \right] ^{-\frac{\beta - \gamma}{\alpha}},
\]

where \( \gamma \) measures the steepness of the core profile, \( \beta \) the outer profile, \( \alpha \) the sharpness of the transition between the two regions and \( I_b \) is the scale factor.
with the centre of the UV isophotes.

Apart from a change in accretion rate, changes in accretion efficiency or circumnuclear obscuration are possible explanations. Circumnuclear material could also provide some reprocessing of the hotter radiation expected from an accretion disk and redate at the cooler 10,000-20,000K observed.

Other explanations for the UV flare based on objects such as binaries, supernovae, stellar flares, outside the nucleus would have to be aligned with our line of sight to such a precision as to make these unlikely candidates. Furthermore, if such were the case, many more\(^3\) such flares should be observed out of alignment with the nucleus in other galaxies and no other such candidates have been seen.

The conclusion must remain therefore that whatever the physical mechanism driving the flare, it is be located within ~1pc of the dynamical centre of the galaxy and is most likely associated with accretion around the central black hole.

\(^3\)The fraction of space within a sphere of radius R which appears within a projected distance r of the centre is \(\sim \frac{3r^2}{2R^2}\) so that we only expect \(\sim 6\%\) (for R=5pc) and \(\sim 1.5\%\) (R=10pc) of the local space to lie within 1pc of the line of sight. We should therefore expect at least 16 the number of off-centre sources in other galaxies if the flaring source was within 5pc of the centre (and \(\sim 66\) sources if within 10pc.) These estimates assume a constant stellar density function within the volume and would be slightly lowered for a steep density function. However the gradients are not so strong as to invalidate the point that if the flare source was associated with the nuclear star cluster rather than the nucleus itself then we would expect one to two orders of magnitude more flares to be seen off-centre than coincident with the nucleus.
5.10.6 The second 1991 UV peak

One proposed explanation for the structure in the 1991 FOC image (Fig 5.1) was that the second UV peak could be due to dusty material surrounding the nucleus providing a working surface for a nuclear jet (Renzini 2001). There is evidence for a dusty torus around the nucleus of NGC 4552 (Carollo et al. 1997). However, if such a jet is also responsible for the radio emission then it does not appear to be aligned with the centres of the two UV peaks. Note that in Fig 5.27 the offset of the UV peaks from the centre of the optical countours (X) is within the errors of the uncertainty in the background subtraction (Cappellari et al. 1999). The separation of the UV peaks in the 1991 observation (\(\sim 0.15''\)) corresponds to \(\sim 45\) light years. If the second peak is any form of light echo of emission from the central source, it is therefore most unlikely to be directly related to the process causing the central emission peak.

Similarly if the emission is due to disruption ejecta interacting with the ISM (e.g. Khokhlov & Melia 1996), then assuming a maximum ejecta escape speed of \(\sim 10^4\text{km.s}^{-1}\) (Rees 1988) it would be due to an ejection event \(\sim 1400\) years ago. Another difficulty for the ejecta model is that the emission from ISM interaction is likely to be more extended than the point source \((r<2\text{pc})\) detected in 1991.

Both of these scenarios seem unlikely especially given the similar total brightness and closely related timescales for the central and off-centre sources.

This spot shows some similarity with a UV 'hotspot' observed close to the
CHAPTER 5. THE SOFT X-RAY VARIABILITY OF NGC 4552

Figure 5.27: The 1991 HST/FOC UV observation (left) from Cappellari et al. (1999) Fig. 1 showing double peaked emission from the nucleus. The 5 GHz (6 cm) VLBA map (right) of NGC 4552 (epoch 2000) from Nagar et al. (2002). Note the different scales, with the entire radio structure fitting within the central contour on the UV image. The radio contours are integer powers of \(\sqrt{2}\), multiplied by the \(\sim 2\sigma\) noise level of 0.8 mJy for NGC 4552. The peak flux-density is 93.8 mJy/beam. The distance between the central peak and the knot to the left (5mas) corresponds to a distance of 0.4pc \((\sim 1.3\ \text{lyr})\) at the Virgo cluster distance of 16.8Mpc. This radio map was taken \(\sim 7\) years after the 1993 HST/FOC image showing the brightest UV peak.

centre of the Seyfert 2 galaxy Mrk 477 (Kishimoto et al. 2002). The central \(\sim 2''\) of the galaxy has shown variability in the UV by a factor of \(\sim 2\) (Kinney et al. 1981, de Robertis 1987) but since the hotspot only contributes \(\sim 10\%\) of the UV radiation in this region, it would have to vary by an order of magnitude to explain the overall UV flux variation. Alternatively, the hotspot may be a glimpse through the edge of obscuring material around a central ionising source which would not require the same level of variability in the emission. The authors comment that Sy 2 do not usually show UV/opt variability but the UV variability by a factor of 2 previously observed for this
5.10.7 Future prospects

To date the UV flare seen at the centre of NGC 4552 is the only such example of the phenomenon. Assuming that the flare seen from 1991-1996 was not an isolated event but part of ongoing accretion processes in the nucleus of the galaxy, it would be interesting to acquire further high resolution imaging (and if possible polarimetry) of the nuclear region. To detect more such UV flares, high-resolution UV imagers will need to be flown and their results compared with the existing HST/FOC and HST/STIS archives. At present HST/STIS is the only operating UV imager available and this does not have a polarimetric imaging mode.

At X-ray energies, the proposed 40ks AO-2 XMM-Newton observation of NGC 4552 and any future Chandra follow-ups will give a better indication of the variability of the nucleus, and an examination of changes in the X-ray spectra at these wavelengths may 'throw light' on the physical processes underlying the 1993 UV flare event.
6.1 Introduction

The large amplitude soft X-ray flares seen in galactic nuclei (§2.4) have allowed an initial estimate of the rate at which stellar disruption flares may occur and of the evolution of the flares themselves. The limits from observation (e.g. Komossa 2001, Donley et al. 2002) are broadly consistent with the theories presented in §2.2 but do not as yet strongly constrain the models.

The lack of any current soft X-ray survey mission means that further examination of the ROSAT All-Sky Survey represents the best opportunity for constraining stellar disruption models. The limited follow up observations mean that in most cases flare candidates are identified as a single 'high state' (with AGN-level luminosities) either followed or preceded by a 'low state' with luminosities typical of normal galaxies.
6.2 ROSAT All-Sky Survey

The ROSAT All-Sky Survey (RASS) (Voges et al. 1999) remains the only large coverage survey of the whole sky at soft X-ray energies (Fig 6.1). RASS went some 20 times deeper than the Einstein Medium Sensitivity Survey (Gioia et al. 1990) and detected sources over the whole sky down to 0.05 counts/sec. Its sensitivity down to 0.1 keV also made it ideal for finding the steep spectrum soft sources which have gone undetected by other observatories.
X-ray surveys using ROSAT to search for soft sources (Vaughan et al. 2001, Grupe, Thomas & Beuermann 2001) have been found an efficient way of detecting ultrasoft AGN (steep X-ray spectra with little emission above 0.5keV) which in many cases are optically classified as narrow line Sy 1s based on their strong, narrow permitted emission lines and ratio of \([\text{[OIII]}]/H\beta \leq 3\). These surveys also detect the soft X-ray flares described in §2.4 although most of the ultrasoft sources do not display the very strong variability of the flares. In most cases (see the lower part of Table 2.1) the optical spectra of the flares do not show evidence of emission lines indicating either a different physical mechanism or viewing geometry, or that the emission lines had disappeared by the time of the optical follow-up (or were absent in a pre-flare spectrum).

### 6.3 Existing X-ray surveys of disruption flares

As discussed in §2.3 the predictions of the stellar disruption rate are based on black hole mass and stellar densities in the nuclei of galaxies (e.g. Rees 1988, Magorrian & Tremaine 1999).

Using such models Sembay & West (1993) predicted (based on Eqn 2.8) the rate of flaring expected to be detectable during the RASS. They were principally concerned with how common super massive BHs were and depending on the assumed mass distribution they predicted between \(\sim 10^2\) and \(\sim 10^3\) flares to be observable during RASS.

Recently, Donley et al. (2002) have calculated a flaring rate in the local
CHAPTER 6. FLARES IN X-RAY SURVEYS

\(z \leq 0.091\) universe of \(\sim 9.1 \times 10^{-6}/\text{year}\) for normal galaxies and \(\sim 8.5 \times 10^{-4}/\text{year}\) for active galaxies based on the expected number of galaxies not detected in a flaring state during RASS relative to subsequent pointed PSPC observations. These rates are broadly consistent with those predicted in §2.3. The sky coverage for this survey was \(\sim 9\%\) based on the requirement that \(|b| > 30\°\) and a PSPC pointing be within 50’ of the axis. Donley et al. did not find any other variability above a factor 20 in their survey meaning that the objects listed in Table 2.1 remain the only X-ray flare candidates.

The total PSPC pointed archive only covers 18% of the sky (and the HRI, 2 %) so there are no good soft X-ray comparison observations for the majority of the RASS sources. Some fraction of these will be expected to be normal galaxies in outburst. Comparisons against previous observations (Einstein, EXOSAT) do not in general have either the spatial resolution or the sensitivity to provide detections (or useful upper limits). For faint sources, ASCA is almost useless due to its poor spatial resolution leading to higher source confusion and limiting sensitivity. The current sky coverage by Chandra, XMM, RXTE is running at few percent per year and does not contribute greatly to all-sky statistics. The major statistical contribution from these narrow field, high resolution observations will be in deep pencil beam surveys (although as discussed in §6.6.6 there are limits to these imposed by the nature of the flares themselves). Of course, high spatial resolution observations of these sources will provide data on the regions responsible for the flare (central pc, binaries in bulge/halo.)

The only search to date for variability in a catalogue of known galaxy posi-
tions is that by Komossa (Komossa 2001). This survey examined the 136 of 486 galaxies in the Ho, Filippenko & Sargent (1995) sample\(^1\) with at least two ROSAT observations and found no evidence for source flaring. However, the small sample size places an upper limit to the flaring rate which is more than two orders of magnitude higher than the upper end of the predicted rates discussed in §2.3.

One remaining area which has not been fully explored is to search for normal galaxies which were not detected during RASS but were detected marginally in a pointed observation. This class of 'significant' non-detections in the RASS will give another indication of variability (even if that is much smaller than the factor \(\sim 100\) for the large amplitude flares) and identify any candidates which may form the fainter end of the flare distribution (or be late decay tails from flares.) The aim of the following analysis is to determine the occurrence of significant non-detections during RASS and determine if they are associated with normal galaxies.

### 6.4 Selecting an optical sample

In order to determine the frequency of flares from stellar disruption and the nature of the flares themselves, a much larger population of sources is needed than the \(\sim 500\) galaxies in the Ho, Filippenko & Sargent (1995) sample. The

\(^1\)Ho, Filippenko & Sargent (1995) constructed a statistically complete sample of 486 galaxies with \(B_T \leq 12.5\) mag and \(\delta > 0^\circ\) in order to search for low-luminosity AGN in the nearby universe.
CH APTER 6. FLARES IN X-R AY SURVEYS

LEDA\textsuperscript{2} database (Paturel et al. 1997) was selected based on the fact that it was an all-sky collection containing \( \sim 10^6 \) galaxies. It is a collation of available data from the literature with one of its major constituents being the Principal Galaxy Catalog (Paturel 1989). Astrophysical parameters and their uncertainties are treated in a consistent way to minimise the systematic effects between catalogues.

For the present work, the sample includes all galaxies with measured redshifts and with \(|b| > 20^\circ\) which yielded 124,684 sources. The \(|b| > 20^\circ\) condition excludes the bulk of the galactic plane where, in general, there is much more interstellar absorption and the probability of source confusion with galactic objects is higher. As a further condition a measured B magnitude is required to allow correlation between optical and X-ray luminosity. The resulting selections are summarised in Table 6.1. These 116,711 galaxies are used as the source positions in the subsequent X-ray analysis.

| LEDA \(|b| > 20^\circ\) and \(V > 0\) and \(B_{\text{mag}} > 0\) | 939,217 |
|-----------------------------------------------------------|---------|
| E/E-SO                                                    | 4859    |
| SO/SO-a                                                   | 6297    |
| SB                                                        | 5525    |
| S                                                         | 16754   |
| Irr                                                       | 1075    |
| No type                                                   | 82201   |

Table 6.1: The selection from the full LEDA catalogue of galaxies with a measured redshift (\(Z>0\) i.e. recession velocity \(V>0\)) and also having a measured B magnitude. The distribution of morphological types for this selection is shown.

\textsuperscript{2}http://leda.univ-lyon1.fr
CHAPTER 6. FLARES IN X-RAY SURVEYS

6.5 Flare Models

The flare models outlined in §2.2 are expected to be visible in the soft X-ray and UV at peak output and to follow a decay as $t^{-5/3}$ (assuming that light output follows mass fall back). Following the model of NGC 5905 (Table 6.2) the luminosity will drop by a factor $\sim 35$ to $10^{41}\text{erg.s}^{-1}$ over the first two years. From this it is reasonable to expect a flaring source still to be detectable within a year of disruption.

The existing studies of the flaring rate during the RASS have assumed that if a flare had happened during 1990 that it would be visible in RASS. For nearby sources this is reasonable but for sources significantly affected by galactic absorption or redshifted, the time during which the flare will be observable above a detection threshold shortens with respect to the nearby objects.

<table>
<thead>
<tr>
<th>$t$ (yrs)</th>
<th>$L_X (10^{42}\text{erg.s}^{-1})$</th>
</tr>
</thead>
<tbody>
<tr>
<td>0.12</td>
<td>10</td>
</tr>
<tr>
<td>0.24</td>
<td>3.16</td>
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<td>0.00316</td>
</tr>
<tr>
<td>30.64</td>
<td>0.001</td>
</tr>
</tbody>
</table>

Table 6.2: The distribution of flare luminosities with time for the NGC 5905 model of Li, Narayan & Menou 2002 showing the doubling of the time intervals in which each luminosity interval would be detected. The disruption time, $t_d$, occurs in the range $t=0.15-0.23\text{years}$.
Table 6.3: These flares were bright during RASS and subsequently fainter. For a change in flux by $F/F_0$ over ΔT years since the RASS (when $t=0$), the disruption of the star would have occurred at time $t_d$ if the subsequent drop in output is to follow the $(t - t_d)^{-5/3}$ decay law. RX J1420.4+5334 was observed in a PSPC pointing before the RASS was carried out and is the best constraint to date on the onset of the flare.

Following the analysis of Li, Narayan & Menou (2002) of the light curve of NGC 5905, Table 6.3 shows the calculated disruption times based on the subsequent decay behaviour observed. In most cases (apart from IC 3599) this is simply a fit of the decay law to two data points but it nevertheless allows a calculation of the expected disruption time relative to the peak brightness observed. The variations in post-disruption interval before the observation may be partly explained by differences in BH mass and stellar orbital parameters but it is clear that simple fitting of the $t^{-5/3}$ decay law may not be appropriate in all circumstances.

Despite these concerns, throughout the following analysis the standard model of a flare (i.e. NCG 5905) is used with the assumption that a flare will remain detectable for $\sim$1 year.
CHAPTER 6. FLARES IN X-RAY SURVEYS

6.6 RASS Data Reduction

6.6.1 Organisation of the RASS data Archive

The RASS data has been organised into 6° × 6° fields (see Fig 6.2) over the whole sky with sufficient fields in each declination strip to achieve an overlap of not less than 0.23°. The optical positions from the LEDA catalogue were separated into tables for each declination strip to produce manageable list sizes (Table 6.3). The density of galaxies per field depends on the richness of the region and the level of data available in the literature for inclusion in the LEDA catalogue. The total exposure time in a field depends on the

![Field structure of the 1378 data sets in the RASS Archive organised in declination strips of 6.4° each.](image)

Figure 6.2: Field structure of the 1378 data sets in the RASS Archive organised in declination strips of 6.4° each. The ecliptic poles are marked within the oval regions. RASS datasets are numbered as rs93aabb where aa is the declination zone number and bb is the field number within that zone. Each field overlaps with its neighbouring fields by at least 0.23°
number of times the satellite scanned that region with the typical exposure time being ~400 seconds (corresponding to two days of strip scans each taking 90mins). The fraction of the sky with exposure time $\geq 100$s is $\sim 97\%$ (Voges et al. 1999). Due to automatic instrument switch off during passage through the South Atlantic Anomaly and other auroral radiation triggers, some patches of the sky have lower (or zero) exposure (see Fig 5.3). These observation gaps were filled in with subsequent pointed observations in early 1991 and the data included in the combined RASS fields. The north and south ecliptic poles were each observed every 90 minutes during the survey and have total integration times over 20,000 seconds (Fig 6.3).

![Figure 6.3: Integrated and differential plots of the sky fractions exposed against RASS exposure time. The histogram shows the fraction of the sky covered with exposure time $t_0$. The dashed and dotted lines show the fraction of the sky with exposure time less than $t_0$ or greater than $t_0$, respectively (from Voges et al. 1999 Fig 2.)](image-url)
### Table 6.4: Distribution of LEDA objects with RASS field declination.

The LEDA/Row and LEDA/Field values are only an estimate due to the field overlap of $>0.23^\circ$. Problems were encountered during the original RASS analysis due to there being $>9999$ sources in the tables for rows 17, 18, 22 and 24 which led to the COMPUTE/UPPER.LIMITS bug being discovered as discussed in §6.6.7.

<table>
<thead>
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<th>RASS Row</th>
<th>RASS Fields</th>
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<td>4062</td>
<td>63</td>
</tr>
<tr>
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<td>64</td>
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<td>3780</td>
<td>59</td>
</tr>
<tr>
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<td>64</td>
<td>-2.8125</td>
<td>11923</td>
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</tr>
<tr>
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</tr>
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<td>58</td>
<td>-30.9375</td>
<td>17494</td>
<td>302</td>
</tr>
<tr>
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<td>55</td>
<td>-36.5625</td>
<td>6706</td>
<td>122</td>
</tr>
<tr>
<td>24</td>
<td>52</td>
<td>-42.1875</td>
<td>14846</td>
<td>286</td>
</tr>
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<td>48</td>
<td>-47.8125</td>
<td>8089</td>
<td>169</td>
</tr>
<tr>
<td>26</td>
<td>43</td>
<td>-53.4375</td>
<td>2102</td>
<td>49</td>
</tr>
<tr>
<td>27</td>
<td>39</td>
<td>-59.0625</td>
<td>2009</td>
<td>52</td>
</tr>
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<td>28</td>
<td>33</td>
<td>-64.6875</td>
<td>1582</td>
<td>48</td>
</tr>
<tr>
<td>29</td>
<td>28</td>
<td>-70.3125</td>
<td>554</td>
<td>20</td>
</tr>
<tr>
<td>30</td>
<td>22</td>
<td>-75.9375</td>
<td>332</td>
<td>15</td>
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<td>31</td>
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<tr>
<td>32</td>
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<td>-87.1875</td>
<td>112</td>
<td>11</td>
</tr>
<tr>
<td>33</td>
<td>1</td>
<td>-90.0</td>
<td>10</td>
<td>10</td>
</tr>
</tbody>
</table>
6.6.2 Computing requirements

The uncompressed data size of the RASS after processing is approximately 107GB. It was not feasible to store this data uncompressed (since the ROSAT Pointed Archive was also stored comprising several times this amount.) Instead the data is stored compressed, expanded set by set during analysis and then recompressed. The speed of compression/decompression as well as the compression ratio is important since the data analysis pipeline is designed to be able to re-analyse the RASS archive with any input catalogue of positions. Since the data comprises photon event lists and engineering logs, the total size of a given RASS field is mainly driven by the number of bright sources it contains and also the total exposure time. More bright sources in a field lead simply to longer photon lists whereas longer exposures also have more engineering data.

Since the data had to be decompressed/compressed during the analysis pipeline, execution time as well as compression factors had to be considered. In order to determine which was the optimum compression scheme from the standard packages available, I chose some typical data set sizes and checked the performance of the standard compression routines on these data. As can be seen from Table 6.5, the optimum compression is achieved for the bzip2 though at a significant speed penalty. Disk IO time is not the dominant factor in the compression/decompression times so even though caching has some impact on these timing exercises compared to real pipelines, the relative performance between compression schemes will remain.
Table 6.5: Compression ratios on sample small (~7MB), medium (~10.7MB), large (~18MB), and maximum (~123MB) sized RASS data sets (NB These sizes refer to the final ‘gzip -1’ compressed size of the data.) The resulting sizes for each compression scheme are shown as a percentage of the original raw size. Execution times in seconds (in parentheses) are given for an unloaded 1.1GHz Athlon (133MHz FSB, with 512MB PC133 SDRAM) with the data stored on a 7200rpm 80GB EIDE disk. The decompression times (in seconds) in the bottom 3 rows vary by 10-20% around the values shown depending on compression ratio for gzip and bzip2.

The distribution of the 1378 data set sizes is shown in Figure 6.4. Since the data size of the majority of the RASS fields corresponds to the 'Medium' category, it was decided to use 'gzip -1' which gives the best speed/compression performance for Medium/Small data sets and doesn’t degrade too badly for the larger sizes.

### 6.6.3 Source detection & upper limit calculation

The standard EXSAS commands for reading and converting the Rationalised Data Format (RDF) FITS files into MIDAS format tables were used and stan-
standard source detection pipelines used. A source was classified as a 'Detection' if it had a Maximum Likelihood (ML) $\geq 10$ (approximately 4 Gaussian $\sigma$). If ML $< 10$ then the 95% confidence upper limit count rate was calculated at the optical position of the galaxy and the source was classified as an 'Upper Limit'. These classifications will be used throughout the remainder of this work.

The upper limit calculation involves finding the highest local excess above background within a search radius of $5 \times$ (FWHM of the PSF) of the optical coordinate so some offset in the X-ray position typical results. The full pipeline listing is given in Appendix A with some annotation.
6.6.4 Exposure time calculation

In the APR01 EXSAS release, exposure time is not determined automatically for RASS data since it isn't stored as a single keyword value in the original data files (the method used for pointed observations). Instead to calculate the effective exposure for each source, the average of the exposure map values within a square of $\sim 100'' \times 100''$ centred on the source is used. For all areas except the ecliptic poles (see Fig 6.8b, top right) the exposure time does not change very much across this scale. The exact area used for exposure calculation varies slightly from 25, 30 to 36 pixels since the exact number of $45'' \times 45''$ pixels included changes depending on the source position within a pixel.

6.6.5 Field overlaps

In order to allow detection of structures near the RASS field boundaries (see Fig 6.5), all fields overlap by at least $0.23^\circ$. Applying the source detection routines to adjacent fields produces multiple detections (or upper limits) for sources within the overlaps. If sources are too close to the field edge for an exposure time extraction region to be measured then the position is flagged as un-determined and are not used for subsequent analysis. The remaining detections within the overlap area with good extraction flags always produce the same values in each field and only one copy is used for the subsequent database. For upper limit calculations there is some variation due to the possibility the highest background fluctuation within the search area may be
in a neighboring field. This leads to a lower upper limit being given in the clipped field. In the case of two (or more) upper limits for a particular source the highest is chosen as the most accurate upper limit determination.

6.6.6 Luminosity calculations

The RASS count rates were converted to fluxes using a 0.05keV black body model with an absorbing column fixed at the galactic value. This is the typical model used in the literature discussed in §2.4 and to enable comparison with other results this model has been followed. For each galaxy position, the galactic absorbing column from Dickey & Lockman (1990) was read using the 'nh' FTOOLS\(^3\). If there is further absorption in the host galaxy then the original source flux would be higher.

Due to the steepness of the spectra being considered here, account must be taken of the change in flux due to redshifting of the spectral shape (K correction, see Fig 6.5 and 6.6). Here I do not use magnitude based K-corrections (Humason, Mayall & Sandage 1956, Oke & Sandage 1968) but express the flux correction equations in terms of energy.

Initially, the effects of change in energy range and spectral shape due to the redshift are considered. For an object at redshift \(z\), the flux density, \(f_E\) emitted at energy \(E_2\) will be observed at energy \(E_1 = E_2/(1 + z)\) with a corresponding flux,

\[
f_O(E_1) = \frac{f_E(E_2)}{1 + z}.
\]

\(^3\)http://heasarc.gsfc.nasa.gov/doc/dos/software/ftools/
Figure 6.5: Black body spectra for kT=0.05keV at redshifts of 0, 0.3 and 1.

Relating the flux emitted at $E_2$ to that emitted at $E_1$ for a spectrum of the form, $f(E) \propto E^{-\Gamma}$ gives,

$$f_E(E_2) = \frac{f_E(E_1)}{(1+z)^{\Gamma}}.$$  \hspace{1cm} (6.2)

Substituting Eqn 6.2 into Eqn 6.1 gives

$$f_O(E_1) = \frac{f_E(E_1)}{(1+z)^{1+\Gamma}}.$$ \hspace{1cm} (6.3)

For $q_0 = 0$, the luminosity distance $d_L = cz(1 + z/2)/H_0$, and therefore the emitted luminosity in a given band will be related to the observed flux in the same band by,

$$L_X = F_X(N_H) 4\pi \left(\frac{c.z (1 + z/2)}{H_0}\right)^2 (1 + z)^{\Gamma+1}.$$ \hspace{1cm} (6.4)
where $F_X$ is the flux corrected for absorption. For spectra with $kT_{bb}=0.05\text{keV}$ the slope of the high energy tail is approximately modelled by a powerlaw with $E^{-4}$. Therefore, for the calculation of luminosities from each of the count rates, Eqn 6.4 was used with the absorption corrected flux (calculated using PIMMS for $kT_{bb}=0.05\text{keV}$, $N_H$ (galactic)), the galaxy redshift and $\Gamma=4$. This gives fluxes within 5% of the thermal model in the energy range $0.1$-$1.0\text{keV}$ and is much faster to calculate (using Eqn. 6.4) for the large number of sources in the catalogue.

The steepness of the flare spectra means that the observed flux is especially sensitive to redshift. As shown in Fig 6.5 the peak energy moves closer to the $\sim 0.1\text{keV}$ edge of the ROSAT PSPC sensitivity (see §3.4.3) with increasing redshift and it is also moved into the region heavily absorbed by material in our own galaxy.

The redshifting of the spectrum coupled with our local absorbing column means that there is an effective cut-off on the distance to which we can see flares.

For a source detection threshold of $0.01\text{cts/sec}$ in the ROSAT PSPC (Fig 6.6), a flare at $10^{42}\text{erg.s}^{-1}$ will be undetectable at $z=0.1$ even at low $N_H$ and the expected peak luminosities ($\sim 10^{44}\text{erg.s}^{-1}$) will be undetectable at $z=0.1$ for $N_H > 10^{21}\text{cm}^{-2}$.
Figure 6.6: The minimum luminosity required for a black body (kT=0.05keV) to be detected at a ROSAT PSPC count rate $\geq 0.01$cts/sec. The curves reading from the bottom correspond to $N_H = 3 \times 10^{19}$, $10^{20}$, $5 \times 10^{20}$, and $10^{21}$cm$^{-2}$. The curvature at higher redshift is due mainly to the contribution from k corrections.

6.6.7 Problems during RASS data processing.

No photons in the source

For some upper limit calculations the algorithm fails to find any peak in the search area (chosen to have radius $5 \times$FWHM). This leads to a failure to calculate a source upper limit. Nevertheless a background count level is available. From that I have examined the relationship between background and upper limit for all of the correctly calculated upper limits (See Fig 6.7). The lower edge of the of the counts distribution for ML=10 (detected) sources gives a reasonable estimate for a conservative upper limit. I have fit this approximately with Equation 5.1 and this is used to generate an upper limit
when none was available from automatic detection.

\[ UPR\_LIMIT\_COUNTS = 3 + (12 \times BGND) \]  \hspace{1cm} (6.5)

This gives a more realistic estimate for the upper limit than a fixed mean value due to the strong dependence of the upper limit on background count level above 0.1 counts. Rather than risk underestimating the upper limit a floor of 3 counts is set. This is approximately twice the lowest upper limit in the database. For 77 sources the upper limit algorithm converges but does not find any photons to include in the source. This is mainly due to short exposure time with 75 of the 77 having exposures below 40 seconds. For these 77 sources Equation 5.1 is also used to calculate an upper limit on the counts from the background map. Sources with the calculated upper limit are given a fixed vignetting factor of 1.5, the modal value for the correctly calculated fraction of the galaxies.

**Bug in COMPUTE/UPPER\_LIMITS routine**

During the processing of the data a bug was discovered in the EXSAS routine COMPUTE/UPPER\_LIMITS in which tables with more than 9999 source positions were not properly handled (see Table 6.3). Sources 1 to 9999 are handled correctly but source 10000 is skipped and instead the upper limit for source 10001 is produced in the output file. This substitution of source \((n+1)\) continues to the end of the list. The second last entry is again correct since the routine cannot read past the end of the input source list and instead
Figure 6.7: The effect of the background count level on the source detection algorithm. Detected sources (ML>10) are plotted in blue and upper limits in red. The curve in green was fit to approximately follow the lower edge of the source detections (blue) and is used to calculate conservative upper limits where there is no automatic solution available.
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uses the last entry.

The form of the problem is shown in the following summary table output from COMPUTE/UPPER_LIMITS:
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Sequence INDEX IND_INP ..... etc ..... 
--- --- --- --- --- --- --- ---- ---
1 1 1 Good

9999 9999 9999 Good
10000 10001 10001 <- 10000 skipped
10001 10002 10002

n n+1 n+1 All rows refer to the (n+1) source in the input list.

24999 25000 25000 Last two elements are identical
25000 25000 25000 Last element is correct again.

I reported this problem to Dr. Rainer Gruber at MPE in Garching who confirmed the bug and very kindly provided a bug fix for the compupp routine. The recompiled EXSAS modules were inserted and the data analysis pipelines were repeated without problem.

Too many sources in a RASS field

During the analysis or the RASS data, seven of the 1378 datasets were highlighted with problems. These problem sets are listed in Table 6.6 and shown in Figs 6.8a and 6.8b. Apart from the detected supernova remnants (SNR positions and identifications confirmed from Green (2001)), the bright diffuse emission detected at the North and South Ecliptic Poles (NEP/SEP) created problems. The large exposure times at both the NEP and SEP due to the overlap of the RASS scan strips means that these ~2° diameter regions are seeing down to the soft X-ray background (mainly due to the hot ISM in our own galaxy (McCammon & Sanders 1990)). The EXSAS source analysis
software is optimised for detecting point sources against a uniform background. In regions of bright diffuse emission, the algorithms tend to detect many 'point sources' along structures within the diffuse emission. This leads to large detection lists containing significant false detections. Due to a limitation built into the EXSAS analysis routines, the Maximum Likelihood detection routines cannot handle tables of more than 1000 candidate sources. To continue the analysis past this point, the input list would have to be 'hand-weeded' of false detections to allow the normal detection parameters to be calculated for the un-confused parts of the image.

<table>
<thead>
<tr>
<th>RASS Field</th>
<th>RA</th>
<th>Dec</th>
<th>Sources</th>
</tr>
</thead>
<tbody>
<tr>
<td>930521</td>
<td>263.5714</td>
<td>67.5000</td>
<td>North Ecliptic Pole</td>
</tr>
<tr>
<td>930522</td>
<td>276.4286</td>
<td>67.5000</td>
<td>North Ecliptic Pole</td>
</tr>
<tr>
<td>931148</td>
<td>310.9091</td>
<td>33.7500</td>
<td>Cygnus Loop SNR</td>
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<tr>
<td>931251</td>
<td>313.4483</td>
<td>28.1250</td>
<td>Cygnus Loop SNR</td>
</tr>
<tr>
<td>932517</td>
<td>123.7500</td>
<td>-45.0000</td>
<td>Puppis A SNR, Vela SNR</td>
</tr>
<tr>
<td>932518</td>
<td>131.2500</td>
<td>-45.0000</td>
<td>Vela SNR</td>
</tr>
<tr>
<td>932907</td>
<td>83.5714</td>
<td>-67.5000</td>
<td>LMC, South Ecliptic Pole</td>
</tr>
</tbody>
</table>

Table 6.6: RASS fields with too many sources to be processed with the standard analysis pipeline. The SNR identifications were confirmed from Green (2001). The North and South Ecliptic Poles have total exposure times of over 20ksec where the RASS scan strips overlap. The presence of the Large Magellanic Cloud sources in the 932907 image (as well as the SEP) caused problems.
A scheme to avoid this table size limit has not yet been implemented using the original pipeline on these frames. Instead the following lines associated with the DETECT/SOURCES command are omitted:

```
......
detect/sources image1 events ? nodispl 10 ? ? bacmpi.bdf
trans/coord solst solst_det2k 2000
$mv solst.tbl solst_det.tbl
......
```

This means that only the upper limits at the catalogue positions are calculated for these frames. If the source has ML>10 then it is classified as a detection and treated in a similar manner to the detections by DETECT/SOURCES from all other frames.

A further problem with the extended regions such as SNRs is that the 'background' determination during source detection is significantly higher leading to less useful upper limits. The high background portion of the Vela SNR (Fig 6.8a, lower left) is strongly contributed to by these bright extended sources.
Figure 6.8a: The problem images from the RASS. Each false colour image is 6.4° × 6.4° square with black/blue representing fewest counts/pixel, through red to yellow/white for maximum counts/pixel. Note the sky overlap between adjacent data sets. The top pair of images (930522 left, 930521 right) show the North Ecliptic Pole region. The lower pair (932518 left, 932517 right) show the Vela SNR with the smaller bright Puppis A SNR visible top right in 932517.
Figure 6.8b: The lefthand pair of images (931148 top, 931251 bottom) show the Cygnus Loop SNR. The offset in Right Ascension (X-axis) between these two images is due to the different numbers of tiles per strip for these two declinations (see Fig 5.2). The bottom right image (932907) shows the South Ecliptic Pole region (left middle of image) and also the bright LMC sources. It has a slightly different false colour scale to the other six images shown to enhance the visibility of the extended emission in the background despite very bright sources in the LMC area. The exposure map for the 932907 data is shown (top right) to indicate the peak in exposure time (contours shown in kiloseconds) around the SEP.
6.6.8 Database preparation

Once all fields had been processed, the detection and upper limit files for each field were collated into a database with all duplicates removed. An X-ray identification with the LEDA galaxy was accepted if the source was within 1' of the optical position for detections and up to 3' for upper limits (Fig 6.10). The increased radius for upper limits is due to the highest local excess being determined within a search area (see §6.6.3). The resulting catalogue contains 103,272 upper limits and 7143 detections (at ML\(\geq\)10) with the remaining 6296 galaxies of the original selection falling outside the offset cut-offs. The breakdown of these results by morphological type is shown in Table 6.7 with the associated luminosity distributions shown in Fig 6.9. The detections and upper limits for each morphological type were correlated with B magnitude but no systematic relationship was found to present.

<table>
<thead>
<tr>
<th>LEDA Type</th>
<th>Detections</th>
<th>Upper Limits</th>
</tr>
</thead>
<tbody>
<tr>
<td>S</td>
<td>1717</td>
<td>26654</td>
</tr>
<tr>
<td>E</td>
<td>913</td>
<td>3876</td>
</tr>
<tr>
<td>Irr</td>
<td>48</td>
<td>1008</td>
</tr>
<tr>
<td>No type</td>
<td>4465</td>
<td>71734</td>
</tr>
<tr>
<td>Total</td>
<td>7143</td>
<td>103272</td>
</tr>
</tbody>
</table>

Table 6.7: The galaxy detections and upper limits by morphological type.

6.6.9 Comparison with RASS-FSC/RASS-BSC

In order to ensure that the database is consistent with existing catalogues the count rates and exposure times for the detected sources were compared
Figure 6.9: Comparisons of detections and non-detections for each of the major morphological type in the LEDA catalogue: Clockwise from top left: Ellipticals (E/S0), Spirals, Irregulars, and those with no type assigned in LEDA.

with the those RASS-Bright Source Catalogue (RASS-BSC). Positional coincidence within 1’ was accepted as a match with most objects coinciding to better than 10” (Fig 6.11 top). The count rate ratio $\frac{RASS-BSC\ CR}{LEDA\ Detection\ CR}$ (Fig 6.11 bottom) shows a peak as expected around unity but with significant scatter. This is due to the different source detection routines used (in particular in high background areas) and partly due to extra exposure time included in the RASS final data release. The mean ratio of exposure times (Fig 6.12) for the RASS-BSC sources and the present detections was $0.96 \pm 0.08$ (1σ).
6.6.10 Upper limits

The useful luminosity upper limit determined at a single epoch is the number of sources below $10^{42}\text{erg.s}^{-1}$ since this is the typical luminosity above which a flare can be clearly distinguished from non-AGN emission levels.
Figure 6.11: Offset of positions between detections and RASS-BSC matches (top). Only objects within 1' were accepted as matches. The lower figure shows the count rate ratio $\frac{\text{RASS-BSC CR}}{\text{LEDA Detection CR}}$. 
Figure 6.12: The exposure time ratio between the RASS-BSC and the present analysis shows the RASS-BSC data with marginally shorter exposure times.

Of the 103,272 upper limits determined (Fig 6.13) using the $kT_{bb}=0.05\text{keV}$, $N_H(\text{galactic})$ model, the number of sources below $10^{42}$, $10^{41}$ and $10^{40}\text{erg.s}^{-1}$, are 18,860, 3963 and 687 respectively. At the same luminosity cut-offs, 498, 148 and 40 respectively of the 7143 detections ($ML>10$) are seen. This gives a population of 19,358 galaxies with luminosities clearly below the flaring/AGN threshold.
Figure 6.13: Distributions of flux and luminosity for the LEDA sample. Upper limits (N=103,272) are plotted as a solid line and the detections (N=7143) are shown as a dotted line.
6.7 Correlations with ROSAT Pointed detections

To search for further flares, the RASS upper limits for the LEDA galaxies were compared with detections from the pointed ROSAT PSPC observations. For each LEDA galaxy with a RASS upper limit, the archive was searched for a pointed detection by the ROSAT PSPC using the WGA catalog (White, Giommi & Angelini 2000). Of the 103,272 galaxies positions checked, 953 had pointed detections within 1' in the WGA catalog. The ratio \( \frac{\text{WGA Count Rate}}{\text{RASS Upper Limit}} \) was \(<1\) for 886 galaxies and of the 67 with ratios \(>1\), 13 were \(>2\) and 4 were \(>3\). The distributions are shown in Fig 6.14.

<table>
<thead>
<tr>
<th>RA</th>
<th>Dec</th>
<th>LEDA</th>
<th>NED ID</th>
<th>Type</th>
<th>WGA/RASS</th>
</tr>
</thead>
<tbody>
<tr>
<td>05h01m46s</td>
<td>-22°53'23&quot;</td>
<td>75257</td>
<td>IRAS04597-2257</td>
<td>Sy 1</td>
<td>&gt;11.7</td>
</tr>
<tr>
<td>14h53m41s</td>
<td>18°03'54&quot;</td>
<td>84348</td>
<td>-</td>
<td>?</td>
<td>&gt;3.2</td>
</tr>
<tr>
<td>23h02m01s</td>
<td>15°58'02&quot;</td>
<td>70295</td>
<td>NGC 7465</td>
<td>Sy 2</td>
<td>&gt;4.9</td>
</tr>
<tr>
<td>23h51m13s</td>
<td>20°14'19&quot;</td>
<td>72612</td>
<td>MCG+03-60-031</td>
<td>Sy 2</td>
<td>&gt;3.3</td>
</tr>
</tbody>
</table>

Table 6.8: The four LEDA galaxies with RASS upper limits less than 3 times the level detected by WGA. The RASS/WGA ratio is a lower limit since the RASS count rate is an upper limit.

The four galaxies with a variability \(>3\) were correlated with the NASA Extragalactic Database (NED) to search for matches. In each case a match was found within 20" of the X-ray position (Table 6.7). The NED identification also lists two unidentified galaxies from the UK Schmidt plates within 40" of the position of IRAS 04597-2257. LEDA galaxy 84348 (in Abell 1991) is classified as an SO with Rmag=14.67 (Colless et al. 1993).

What is clear is that none of the 953 galaxies with detections in WGA and
Figure 6.14: Distributions of WGA Ratio = $\frac{\text{WGA Count Rate}}{\text{RASS Upper Limit}}$ for sources with WGA Ratio $\leq 1$ (top, N=886) and those with WGA Ratio $> 1$ (bottom, N=67). The four sources with WGA Ratio $> 3$ are shown in Table 6.5. NB the two sources shown in the rightmost bin represent ratios of 4.9 and 11.7.
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RASS upper limits show flaring of the type seen during the RASS. Furthermore, the levels of activity are consistent with those for active galaxies. This sample gives a limit to the flaring rate of $\sim 10^{-3}$/yr if it is assumed that a flare would have been bright during the pointed observation.

6.8 The NLSy1 Population

As mentioned in §6.1 one group of galaxies that appear to have consistently soft X-ray spectra are the Narrow Line Seyfert 1 (NLSy1) galaxies first described by Osterbrock & Pogge (1985). These have narrow permitted lines, strong Fe relative to H\textbeta and weak [OIII]. Recently, a systematic search for NLSy1s has been made (Williams, Pogge & Mathur 2002) in the Sloan Digital Sky Survey early release data (Stoughton et al 2002) using optical criteria (Williams, Pogge and Mathur (2002), henceforth WPM). This search yielded 150 objects and they searched for matches with the Bright and Faint RASS Source Catalogues (RASS-BSC, RASS-FSC). This led to 45 detections (16 BSC, 29 FSC). There is no distinction in redshift between the detected and undetected populations.

In order to assess the X-ray variability of the WPM optically selected sample I have carried out a search for detections in other X-ray archives. Of the 45 sources 7 had pointed ROSAT PSPC observations and 3 Einstein IPC. No detections were found in the EXOSAT or ASCA archives.

The seven sources with available ROSAT pointed observations are all detected and listed in the WGA catalogue (White, Giommi & Angelini 2000).
None of the remaining 38 sources were observed after RASS so their absence for the WGACAT is not due to non-detection. The first source in Table 6.7 (J014644.82-004043.2) was also detected with the Einstein IPC at a count rate of $31\pm3\text{cts/ksec}$. For the power law model of WPM and galactic $N_H$, this is equivalent to $78\pm8\text{cts/ksec}$ for the ROSAT PSPC.

The sources show count rate variability by factors of a few over the few years between the RASS and pointed observation. A similar level of variation is seen between the Einstein IPC observation and the RASS for J014644.82-004043.2. Although less well constrained, the hardness ratios show some variations with significant change from a soft to a hard state in J171207.44+584754.5.
Table 6.9: The seven sources with pointed detections from the Williams’ (2002) sample of 45 RASS detected NLSy1s. The offsets from the optical position are given as is the off-axis angle of the source for the pointed WGA detection. The hardness ratios given here for both RASS and WGA are $HR = \frac{H-S}{H+S}$ where $S=0.1-0.4\text{keV}$ and $H=0.4-2.0\text{keV}$.

1) Columns z, HR and $\Gamma$ are from Williams et al (2002). (2) Calculated from data from the Heasarc archive. (3) Calculated from the database of ROSAT pointed detections of White, Giommi & Angelini (2000).

<table>
<thead>
<tr>
<th>Object</th>
<th>z</th>
<th>HR</th>
<th>$\Gamma$</th>
<th>RASS (ksec)</th>
<th>Offset (ksec)</th>
<th>WGA (ksec)</th>
<th>Offset (ksec)</th>
<th>Off-axis (ksec)</th>
<th>HR</th>
</tr>
</thead>
<tbody>
<tr>
<td>J014644.82-004043.2</td>
<td>0.083</td>
<td>0.00±0.14</td>
<td>2.6±0.2</td>
<td>140±20</td>
<td>0.2'</td>
<td>47±5</td>
<td>0.18'</td>
<td>1.7'</td>
<td>0.20±0.05</td>
</tr>
<tr>
<td>J030417.78+002827.4</td>
<td>0.044</td>
<td>0.46±0.39</td>
<td>3.3±0.8</td>
<td>53±19</td>
<td>0.2'</td>
<td>38±3</td>
<td>0.12'</td>
<td>10.5'</td>
<td>0.76±0.14</td>
</tr>
<tr>
<td>J165658.38+630051.1</td>
<td>0.169</td>
<td>-0.35±0.11</td>
<td>3.0±0.2</td>
<td>38±5</td>
<td>0.05'</td>
<td>48±2</td>
<td>0.09'</td>
<td>28.0</td>
<td>-0.27±0.04</td>
</tr>
<tr>
<td>J170812.29+601512.6</td>
<td>0.145</td>
<td>-0.16±0.13</td>
<td>2.6±0.2</td>
<td>60±10</td>
<td>0.2'</td>
<td>26±2</td>
<td>0.09'</td>
<td>44.6'</td>
<td>-0.30±0.03</td>
</tr>
<tr>
<td>J170956.02+573225.5</td>
<td>0.522</td>
<td>-0.36±0.16</td>
<td>3.0±0.3</td>
<td>28±5</td>
<td>0.03'</td>
<td>39±5</td>
<td>0.51'</td>
<td>43.7'</td>
<td>-0.43±0.06</td>
</tr>
<tr>
<td>J171207.44+584754.5</td>
<td>0.269</td>
<td>-0.34±0.10</td>
<td>2.9±0.2</td>
<td>60±10</td>
<td>0.33'</td>
<td>61±2</td>
<td>0.12'</td>
<td>2.5'</td>
<td>0.30±0.10</td>
</tr>
<tr>
<td>J171829.01+573422.4</td>
<td>0.101</td>
<td>0.16±0.17</td>
<td>2.3±0.3</td>
<td>39±6</td>
<td>0.11'</td>
<td>132±4</td>
<td>0.08'</td>
<td>26.6'</td>
<td>-0.01±0.08</td>
</tr>
</tbody>
</table>
While no indications for large amplitude variability are present here, the main feature of interest with this full Williams' (2002) sample is the non-detection of X-rays from two thirds of the galaxies. Since they have shown that this is uncorrelated with redshift, it must either be due to differences in emission from the nuclei of the galaxies or differences in circumnuclear absorption. This would have to allow optical NLSy1 features to be observed but no X-ray emission. If such a obscuration was present, it could indicate that a significant fraction of stellar disruption flares are going to be obscured by material in the nuclei of their host galaxies. If the obscuration in NL Sy1s is due to an (approximately) steady mass infall regime (thick torus, etc) then this may not be relevant in the case of the transient stream of material fueling disruption flares. If however, the obscuration is not directly associated with the accretion process then the stellar disruption candidates observed by RASS may represent a particular unobscured class of galactic nuclei with perhaps twice that (from the above considerations) being obscured.
6.9 Discussion

The constraint imposed by the non-detection of a flare from $\sim 10^3$ known galaxy positions is an improvement on that by Komossa (2001) by an order of magnitude. The rate upper limit of $\sim 10^{-3}$/yr is still an order of magnitude above the highest rates predicted (Magorrian & Tremaine 1999, Syer & Ulmer 1999). Until multi-epoch X-ray data are available for $\sim 10^5$ galaxies, it will not be possible to strongly challenge the existing theoretical flaring rates in the range $10^{-4}$-$10^{-9}$/yr.

A further yield from the work is a sample of $\sim 19,300$ galaxies which are clearly not emitting above a level of $10^{42}$ erg s$^{-1}$ during the RASS. This is an important determination for comparison with future surveys since not all areas of the RASS coverage will give any useful constraint due to the short exposure time and high $N_H$.

The remainder of the $\sim 10^5$ luminosity upper limits (Fig 6.13 lower) are less useful in correlating with future surveys due to the limits being similar to the expected flare peak luminosities. However, where sources are in the future observed at $10^{43}$ erg s$^{-1}$ and above, some useful comparisons can be made with the RASS data.

6.9.1 Effects of observable flare duration

As discussed in §2.2 and §2.4, the basic model of a stellar disruption flare is expected a soft thermal spectrum with peak luminosity $10^{42}$-$10^{44}$ erg s$^{-1}$ and
a lightcurve decaying as $t^{-\frac{5}{3}}$. This spectral model is sufficient for the present study with the ROSAT instruments due to their poor spectral resolution but it is not expected to be a full description of the underlying processes in the flare.

These uncertainties regarding flare duration have a significant bearing on the calculated flaring rates since the assumed duration of a flare above the detectability threshold is inversely proportional to the calculated flare rate from a given population of detections. In particular if the early-time decay rate is much quicker than $t^{-\frac{5}{3}}$ then the disruption rate will be much higher based on the present sample of flares than the current estimates of $10^{-5}$-$10^{-6}$/yr (Sembay & West 1993, Donley et al. 2002).

The assumption of equal mass fraction per energy interval for the tidal disruption ejecta (which directly gives rise to the $t^{-\frac{5}{3}}$ decay), may not be the correct model or may deviate significantly in some cases as discussed in . If the stellar material is given a narrower distribution of ejection energies then the interval between the onset and completion of fallback for the bulk of material may be much shorter than suggested by Rees' initial assumption.

6.9.2 Confusion with other classes of flares

There are a number of classes of flaring objects which may be confused with transient accretion onto SMBHs. These can be classified according to spectral and luminosity characteristics and their location either in our galaxy or the target galaxy. Novae and cataclysmic variables in particular within our own
CHAPTER 6. FLARES IN X-RAY SURVEYS

galaxy are likely to show similar fluxes. However, with sufficiently well sample light curves and spectra, these can be discriminated. In the present work no estimate has been made of the probability of a galaxy position being confused with a nearby local galactic object and this correction would need to be included in future assessments. Such a procedure would allow more of the galactic disk below $|b| > 20^\circ$ to be included in surveys. At present the exclusion of the galactic plane minimises the effect of confusion within our own galaxy.

6.10 Future X-ray searches.

The only soft X-ray survey mission planned to be launched in the next decade is Lobster-ISS (Fraser et al. 2002). It contains a micro-channel plate (MCP) imager with instantaneous field of view of $162^\circ \times 22.5^\circ$. Over a 90 minute orbit of the ISS it will image almost the whole sky (apart from sun avoidance). The daily sensitivity of $\sim 2 \times 10^{-12} \text{erg.s}^{-1}.\text{cm}^{-2}(0.1-3.5\text{keV})$ will allow it to detect flares with $kT_{eb}=0.05\text{keV}$ and $L_{peak} \approx 10^{44} \text{erg.s}^{-1}$ to a distance of $z \sim 0.2$ in 1 day and $z \sim 1$ in 1 month. As discussed in §6.6.6 the ability to observe flares beyond $z=1$ is dominated by interstellar absorption of the redshifted spectrum.

The relatively poor spatial resolution of 4' FWHM will mean that detailed identification will be required with high resolution pointed observatories such as XMM or Chandra. However, the launch of LOBSTER in 2009 represents the best prospect for research in flares from galactic nuclei in the near future.
Chapter 7

Conclusions

7.1 Ultraviolet flares

In the case of the UV flare in NGC 4552, it seems clear that whatever process is driving the emission is different from the bright soft X-ray flares seen in the RASS. The upper limit of a few times $10^{40}\text{erg.s}^{-1}$ in the 0.1-0.5keV band for any variability places it far below the $10^{42}\text{erg.s}^{-1}$ or higher typical for the RASS flares.

The flat UV spectrum also makes it unlikely that much X-ray emission was associated with the event through the early 1990’s. The possibility remains however, that the UV emission is a reprocessed component of a hotter underlying spectrum as envisaged by Ulmer and collaborators.

It will be interesting to observe the proper motions and evolution of the radio jet structures over the next decade to determine whether the present knots are likely to have their origin around the time of the peak UV emission. If so this would strengthen the evidence for transient accretion being the
CHAPTER 7. CONCLUSIONS

underlying process for the UV flare with the radio jets only becoming active at times of mass infall.

As yet there is no good explanation for the appearance of the off-nuclear peak in the 1991 FOC observation though models involving the nuclear jet appear unlikely due to alignment differences with the radio data.

7.2 X-ray flare frequency

No further evidence of large amplitude flaring was found in the sample of galaxies with RASS upper limits and pointed PSPC detections.

The upper limit to the flaring rate from this work of \(~10^{-3}\)/galaxy/yr improves on the previous galaxy catalogue attempt (Komossa 2001) but \(~10^{-5}\) galaxies with multi-epoch X-ray detections will be needed to begin to seriously challenge the theoretical model.

The 19,300 galaxies which clearly do not contain flares brighter than \(10^{42}\)erg.s\(^{-1}\) form a useful sample to check against in future surveys. The fact that only a small subset (19%) of the whole provide suitable limits, highlights the fact that despite its sensitivity in the soft X-ray region, the short exposure times, sensitivity to the absorption column and strong reddening effects with redshift, make detecting flares in the RASS difficult beyond the very nearby universe.
CHAPTER 7. CONCLUSIONS

7.3 Future Work

HRI resolution recovery

The results of the wobble detection for NGC 4552 show that it is certainly feasible to detect the amplitude of the residual aspect errors for faint sources even when the source counts are weak.

The next stage for investigation is whether there is systematic behaviour across different observations relating to roll angle and positions of the guide stars in the star trackers. The fixed relationship between the wobble axis and the pointings of the star trackers and X-ray imagers ensure that some observation parameters remain constant. This would allow statistics from several observations to be combined and an improved model of the underlying shape of the correction function to be determined. It is clear that the triangular model used to correct the NGC 4552 data is not sufficient to reduce the source extension.

Despite these obstacles it is a goal worth pursuing. The ROSAT HRI performing at close to its nominal resolution will come close to Chandra/ACIS and be able to place better limits on the variability of LMXBs and particularly on the bright \((L_x > 10^{39}\text{erg.s}^{-1})\) intermediate luminosity objects (IXO) which appear to have masses of \(10^2-10^4\ M_\odot\) (Makishima et al. 2000, Strickland et al. 2001). The variability of the existing ROSAT identified IXO candidates (e.g. Colbert & Ptak 2002) compared with Chandra observations will be important in determining the nature and prevalence of these source. Restoring the resolution of the ROSAT HRI images will allow better upper
limits to be placed on newly discovered Chandra IXO candidates.

**Stellar disruption flares**

The appearance of a stellar disruption flare is expected to be dependent on a wide variety of interaction parameters including stellar type, stellar orbital angular momentum, black hole spin and mass, viewing geometry and obscuration within the host galaxy. It is unlikely that the present small sample of X-ray, UV and optical candidates represents the full range of phenomena associated with stellar disruption. The spectral energy distributions of many more flares will need to be observed to determine which phenomena are associated with each section of the interaction parameter space.

The only clear signature of stellar disruption (compared to other forms of transient accretion) is the characteristic mass fall-back decay. To distinguish disruption flares from changes in accretion this process will need to be better modelled to provide observational tests which can clearly isolate this class.

Further development of the relativistic treatment by Diener et al. (1997) and Ivanov et al. (2002) of disruption by a Kerr black hole will also be needed to predict what role black hole spin plays in the energy distribution of the disruption debris with its implications for the bound mass fraction and energy input into the local ISM.

One important extension to the current work on the LEDA catalogue is to determine distances for the ~90% of objects without a spectroscopic redshift and hence calculate luminosities. This would expand the sample of upper
limits and detections up to $10^6$ objects. Similarly inclusion of the Sloan Digital Sky Survey and 2 degree Fields Survey will add significantly to the catalogue of well characterised galaxies.

The other important relationship to establish is the central BH masses for the galaxies in the sample. These can be estimated through the relations being derived for $M_{Bulge}-M_{BH}$ and between $\sigma$, $L$ and $M_{Bulge}$ (Laor 1998, Magorrian et al. 1998, Gebhardt, et al. 2000, Merritt & Ferrarese 2001, McLure & Dunlop 2002). More recently $M_{BH}$ has been correlated with the bulge shape profile (Graham et al. 2002). These continuing improvements in determinations of black hole masses in galactic nuclei and the associated local stellar velocity distributions will allow far more accurate determination of the efficiency of loss cone repopulation processes for different stellar types, and hence the expected disruption rates.

From the observational side, the next step is to acquire high-resolution X-ray spectra of flare candidates. In the short term this seems likely to be achieved with comparison of observations from XMM and Chandra with the soft X-ray archive and in particular with the RASS. This should yield detections at the rate of a few per year. With the projected launch date for LOBSTER/ISS of 2009, the detection of several hundred flares per year during its 3 year lifespan should broaden the sample considerably and allow the first real statistical characterisation of the distribution of flares in galactic nuclei. LOBSTER’s ability to reach the RASS limits within a couple of weeks and, perhaps more importantly, to detect the expected peak flare output within a few days for will allow alerts for multi-wavelength follow-up observations.
Such observations will allow much more accurate determination of temperatures, ionisation states and densities in the accretion structures. Better sampling of the light curves will also strongly constrain the mass fall-back rate and the nature of the accretion processes.

The characteristically steep soft X-ray spectra of the flares seen to date, coupled with the absorption column within our own galaxy, place an effective limit on the redshift (and hence, lookback time) at which detections can be made with current or planned X-ray observatories. Nevertheless, improvements in the theoretical models and observational constraints from the nearby universe will lead to a better understanding of transient accretion, the contribution of stellar disruption to SMBH growth rate and to the injection of energy back into the circumnuclear ISM from unbound ejecta and radiation driven winds.
Appendix A

EXSAS data reduction of RASS data.

Since the exposure calculation is not a standard part of the source detection pipeline in the EXSAS APR01 release, I review the RASS analysis method I have used. This approach was used for conversion of vignetting corrected upper limit counts and detected counts to count rates throughout the thesis. The first two sections give the source detection and upper limit pipelines respectively with some illustrations of the intermediate data products. The final section gives the full MIDAS procedure used.
APPENDIX A. EXSAS DATA REDUCTION OF RASS DATA.

A.1 Source detection pipeline.

Exsas 001> intape/rdf DISK rs93ABCDn00_anc,bas,mex
   [Where ABCD is the RASS frame ID number]

Exsas 002> create/source_detect_image events 10-240 ? 90
   [3rd param is default pointing direction]
   [4th param is binning in sky_pix (0.5")]

Note. Midas Help for create/source_detect_image states that Survey default is 192 but while this was the case for RASS 1 XXXX (check) release data sets the current 6x6 degree field are binned with 90 sky_pix (45") per ima_pix.

Exsas 003> load/ima image1 cuts=-1,5
   [See Fig XXXX.a ]

Exsas 004> crea/bg_image
   [Edit this file to read as follows.]

| ! @(#)creabg_p.epf          1.3 (MPE-Garching) 1/24/00 21:44:02 |
| ! parameter file for CREATE/BG_IMAGE TASK: CREABG |

| MISSION,DETECTOR | ROSAT,PSPC | mission and detector |
| OBS_MODE | SURVEY | pointing or survey |
| NUM_IMAGES | 1 | number of images (1-6) |
| IMAGE_1 | image1.bdf | first image |
| MIN_ML | 10.00 | minimum ML |
| CUT_RADIUS_FWHM | 3.00 | cut radius in FWHM |
| FIT_BIN_SIZE | 16 | fit bin size [imagepix] |
| LDETECT_LIST_1 | lslist1.tbl | first LDET source list |
| MASK_FLAG | F | flag to use mask (T/F) |
| MASK | mask.bdf | mask |
| EXPOSURE_FLAG | F | flag to use exposure |
| EXPOSURE_IMAGE_1 | exposure1.bdf | first exp image |
| BG_IMAGE_1 | bacmp1.bdf | first bg image |

Exsas 005> detect/sources image1 events ? displ 10 bacmp1
   [2nd param is photon event list.]
   [3rd param is mask which defaults to none.]
APPENDIX A. EXSAS DATA REDUCTION OF RASS DATA.

Table : solst.tbl

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<th>INDEX</th>
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Exsas 007> transform/coord solst solst2k 2000

Exsas 008> read/tab solst2k > solst2k.txt

Table : solst2k

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APPENDIX A. EXSAS DATA REDUCTION OF RASS DATA.

37  74.96436  83.08  120.00  4619.99
119  74.35570  67.67  120.00  6279.10

Exsas 009> load exposure

Exsas 010> statistic/ima exposure CURSOR
For source #37 (NGC 4552)

frame: exposure (data = R4)
   plane_no 1 loaded
area [-4545.50,-8325.50,-2295.50,-10575.5] of frame
minimum, maximum:  4.3319e+02  4.5756e+02
at pixel (206,365),(221,374)
mean, standard_deviation:  4.4367e+02  4.5559e+00
3rd + 4th moment:  0.303939  2.73391
total intensity:  299915
median, 1. mode, mode:  4.4324e+02  4.3323e+02  4.4136e+02
total no. of bins, binsize:  256  9.557236e-02
# of pixels used = 676 from 206,349 to 231,374 (in pixels)

For source #119 (M58)

frame: exposure (data = R4)
   plane_no 1 loaded
area [-8055.50,-3015.50,-5805.50,-5265.50] of frame
minimum, maximum:  4.3157e+02  4.5148e+02
at pixel (170,294),(192,315)
mean, standard_deviation:  4.4074e+02  4.2320e+00
3rd + 4th moment:  0.0364451  2.17907
total intensity:  297940
median, 1. mode, mode:  4.4087e+02  4.3161e+02  4.4168e+02
total no. of bins, binsize:  256  7.805236e-02
# of pixels used = 676 from 167,290 to 192,315 (in pixels)
Figure A.1: The standard binned $512 \times 512$ pixel ($6.4^\circ \times 6.4^\circ$) image from a RASS data set (left). The exposure map (right) for the same data with a greyscale indicating exposures of $\sim 390$ s (black) to $\sim 600$ s (white). The streaking in the exposure map arises from the scan direction of the satellite.
APPENDIX A. EXSAS DATA REDUCTION OF RASS DATA.

A.2 Upper limit pipeline

Exsas 011> write/radec uslst 188.68 12.60 1

Exsas 012> comp/upp

[Edit compupp.epf to read as follows:

```
!
! parameter file for COMPUTE/UPPER_LIMITS TASK: COMPUPP
!
MISSION,DETECTOR     ROSAT,PSPC        ! mission and detector
OBS_MODE              SURVEY            ! pointing or survey
INPUT_DATASET         events.tbl        ! name of events table
AMPLITUDE_LOW         10                ! lower amplitude bounds
AMPLITUDE_HIGH        240               ! upper amplitude bounds
CUT_RADIUS            5.0               ! cut radius in FWHM
OFF_AX_MIN            0.0               ! min off-axis angle
OFF_AX_MAX            57.00             ! max off-axis angle
UL_THR                10.00             ! threshold of -ln(p)
CONF_UPPER_LIMIT      95.40             ! conf. upper lim flux
STARTPOSITION         user_list         ! merged_list / user_list
MERGED_LIST           mplst.tbl         ! merged source list
USER_LIST             uslst.tbl         ! user source list
FIT_POSITION          free              ! fit source position
FIT_EXTENT            free              ! fit source extent
FIT_COUNTS            free              ! fit source counts
BG_IMAGE_1             bacmpl.bdf       ! smoothed bg image
UL_LIST_ACCEPT        solst.tbl         ! list of accept. sources
UL_LIST_ALL           ullst.tbl         ! list of all sources
```

Table : ullst.tbl

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<th>EXI_ML</th>
<th>CTS</th>
<th>CERR</th>
<th>RATE</th>
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<td>UL_RATE</td>
<td>XIMA</td>
<td>YMIA</td>
<td>XERR</td>
</tr>
<tr>
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<td>-------</td>
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<td>---------</td>
<td>------</td>
<td>------</td>
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<td>0.00000</td>
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<th>EXT</th>
<th>EXTErr</th>
<th>EXT_ML</th>
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<th>VIG_COR</th>
<th>OFFAX</th>
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<td>300</td>
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<td>0.70178</td>
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<table>
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<tr>
<td>1</td>
<td>60.00</td>
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</tr>
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A.3 MIDAS code listing

This is the EXSAS/MIDAS procedure used on each field to generate the source detections and upper limits. The ancilliary files used by the routine are indexed through the code and given at the end of the section. Where a lines runs over the right hand edge a continuation is indicated by ' . . '

! ------------------------------------------
! LEDA processing script
! ------------------------------------------

! p1 is directory '93abcd' for indexing source tables
! p2 is the uslstXX to link to for the current dir.

DEFINE/LOCAL num_obj/i/1/1 0
DEFINE/LOCAL i/i/1/1 0
DEFINE/LOCAL j/i/1/1 0
DEFINE/LOCAL rdummy/r/1/1 0.0
DEFINE/LOCAL lx_h/r/1/1 0.0
DEFINE/LOCAL lx_l/r/1/1 0.0
DEFINE/LOCAL ly_h/r/1/1 0.0
DEFINE/LOCAL ly_l/r/1/1 0.0
DEFINE/LOCAL expo/r/1/1 0.0

$rm middum*
$gunzip *
intape/rdf DISK r anc,bas,mex
create/source events 10-240 ? 90
$cp ~/leda/creabg.epf creabg.epf
*** See (1)
$cp ~/leda/deteloc.epf deteloc.epf
*** See (2)
detect/local deteloc
detec/bg_image creabg
detect/sources image1 events ? nodispl 10 ? bacmp1.bdf
trans/coord solst solst_det2k 2000
$mv solst.tbl solst_det.tbl
!------------------------------------------

$cp ~/leda/compupp.epf compupp.epf
*** See (3)
$rm uslst.tbl
APPENDIX A. EXSAS DATA REDUCTION OF RASS DATA.

$ln -s ~/leda/{p2} uslst.tbl  
*** See (4)
comp/upper compupp
trans/coord ullst ullst2k 2000
trans/coord solst solst2k 2000

! ---------------------------------------------------------
! [ For each line in solst & ullst determine exposure around
! coord and calculate CR]
! ---------------------------------------------------------

show/tab uslst
num_obj = {outputi(2)}

crea/col ullst :EXPO
crea/col ullst :NH

do i = 1 {num_obj}
! write/out {i}, {ullst,:REMARKS,{i}}
  if M$VALUE(ullst,:REMARKS,{i}) .ne. "no bgr avail" then
  if M$VALUE(ullst,:REMARKS,{i}) .eq. "no solution " then
    ! ie a zero photon source so use a conservative upper
    ! limit for this position based on the background. See
    ! Chapter 5.XXXX for description.
    rdummy = 3.0 + ({ullst,:BGR,{i}} * 12.0)
    ! Determined from fits to CR vs
    write/tab ullst :UL_CTS {i} {rdummy}
    write/tab ullst :VIG_COR {i} 1.5
    ! write/out "New Calc: " {rdummy} {i}
  ! Regenerate missing coord info for 'no solution' sources

$rm temp.coord.tbl
crea/tab temp_coord
crea/col temp_coord :RA_DEG
crea/col temp_coord :DEC_DEG
write/tab temp_coord :RA_DEG 1 {uslst,:RA_DEG,{i}}
write/tab temp_coord :DEC_DEG 1 {uslst,:DEC_DEG,{i}}

transf/radec temp_coord temp_pix image1
write/tab ullst :XIMA {i} {temp_pix,:XIMA,1}
write/tab ullst :YIMA {i} {temp_pix,:YIMA,1}
write/tab ullst :XSKY {i} {temp_pix,:XSKY,1}
write/tab ullst :YSKY {i} {temp_pix,:YSKY,1}
write/tab ullst2k :RADEG {i} {temp_pix,:RA_DEG,1}
write/tab ullst2k :DECDIG {i} {temp_pix,:DEC_DEG,1}

! Beware: this may produce |X_SKY| > 23,000 which is out of the
! ref_ima for subsequent stat_ima, etc tasks.

endif ! 'No Solution' repairs finished

! write/out "Determining Exposure Time"
lx_l = {ullst,:X_SKY,{i}} - 200.0
lx_h = {ullst,:X_SKY,{i}} + 200.0
ly_l = {ullst,:Y_SKY,{i}} - 200.0
ly_h = {ullst,:Y_SKY,{i}} + 200.0

! Exposure is determined from 5x5 or 6x6 pixels around source
! depending on exact rounding of coordinates to image pixels.

if {lx_l} .gt. -23000.0 .and. {lx_h} .lt. 23000.0 then
  if {ly_l} .gt. -23000.0 .and. {ly_h} .lt. 23000.0 then
    stat/ima exposure [{lx_l},{ly_l};{lx_h},{ly_h}]
    expo = {outputr(3)}
    write/tab ullst :EXPO {i} {expo}
    if expo .gt. 0.0 then
      rdummy = {ullst,:UL_CTS,{i}} / {expo}
      write/tab ullst :UL_RATE {i} {rdummy}
    else
      write/tab ullst :UL_RATE {i} 999.9
    endif
  else ! If it passes X range but fails Y range check
    ! Indicate that src too close to edge to
do determine EXPO
    write/tab ullst :EXPO {i} -999.0
    write/tab ullst :UL_RATE {i} -999.0
    write/out 'Out of Bounds'
  endif
else ! If it fails on X range check
  ! Indicate that src too close to edge to determine EXPO
  write/tab ullst :EXPO {i} -999.0
write/tab ullst :UL_RATE {i} -999.0
write/out 'Out of Bounds'
endif ! out of bounds checking

! Calculate nh using
$rm temp_tab.tbl
$cp ~/leda/nh_value.fmt nh_value.fmt
$nh 2000 {ullst2k,:RA_DEG,{i}} {ullst2k,:DEC_DEG,{i}} ...
    ... | tail -n 1 > nh_value.txt
create/tab temp_tab 1 1 nh_value.txt nh_value.fmt
write/tab ullst :NH {i} {temp_tab,:NH,1}

! else
! write/out 'No Bgr ' {i}
endif ! Is 'No bgr avail'
enddo

! Generate_table
! --------------------------------------
$rm upp_out.tbl sol_out.tbl upp_fin.txt upp_out.fits ...
    ...
create/tab upp_out 1 100000
create/col upp_out :ROS_SEQ "" A15 C*15
write/tab upp_out :ROS_SEQ @1..100000 {pi}
copy/tt uslst :LEDA upp_out :LEDA
copy/tt ullst2k :RA_DEG upp_out :RA_DEG
copy/tt ullst2k :DEC_DEG upp_out :DEC_DEG
copy/tt ullst :EXI_ML upp_out :EXI_ML
copy/tt ullst :CTS upp_out :CTS
crea/col upp_out :EXPO
copy/tt ullst :EXPO upp_out :EXPO
copy/tt ullst :RATE upp_out :RATE
!copy/tt ullst :ERATE upp_out :ERATE
copy/tt ullst :UL_CTS upp_out :UL_CTS
copy/tt ullst :UL_RATE upp_out :UL_RATE
copy/tt ullst :BGR upp_out :BGR
copy/tt ullst :NH upp_out :NH
APPENDIX A. EXSAS DATA REDUCTION OF RASS DATA.

copy/tt ullst :VIG_COR upp_out :VIG_COR
copy/tt ullst :REMARKS upp_out :REMARKS
outdisk/fits upp_out.tbl upp_out.fits
$fdump upp_out.fits upp_out.txt clobber=yes pagewidth=256 ...
... columns="*" rows="-"
$source ~/midwork/trim_upp.csh
$cat ~/leda/concat.uppers upp_fin.txt > ~/leda/concat.uppers2
$mv ~/leda/concat.uppers2 ~/leda/concat.uppers

create/tab sol_out
create/col sol_out :ROS_SEQ "" A15 C*15
write/tab sol_out :ROS_SEQ @1..4096 {pi}
copy/tt solst2k :RA_DEG sol_out :RA_DEG
copy/tt solst2k :DEC_DEG sol_out :DEC_DEG
copy/tt solst :RATE sol_out :RATE
copy/tt solst :ERATE sol_out :ERATE
copy/tt solst :EXI_ML sol_out :EXI_ML
copy/tt solst2k :OFFAX sol_out :OFFAX
outdisk/fits sol_out.tbl sol_out.fits
$fdump sol_out.fits sol_out.txt clobber=yes pagewidth=256 ...
... columns="*" rows="-"
$source ~/midwork/trim_sol.csh
 ! End of generate_tab
 ! ----------------------------------------------------------
 ! Cleanup
$rm uslst.tbl ! Sym-link only
$gzip -1 *
 ! ----------------------------------------------------------
 ! End

Note (1) creabg.epf

See file in Section B.1.
### APPENDIX A. EXSAS DATA REDUCTION OF RASS DATA.  

#### Note (2) deteloc.epf

<table>
<thead>
<tr>
<th>Parameter</th>
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<tr>
<td>MISSION,DETECTOR</td>
<td>ROSAT,PSP</td>
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<tr>
<td>OBS_MODE</td>
<td>SURVEY</td>
</tr>
<tr>
<td>NUM_IMAGES</td>
<td>1</td>
</tr>
<tr>
<td>IMAGE_1</td>
<td>image1.bdf</td>
</tr>
<tr>
<td>MIN_ML</td>
<td>1.000000E+01</td>
</tr>
<tr>
<td>MASK_FLAG</td>
<td>F</td>
</tr>
<tr>
<td>MASK</td>
<td>none</td>
</tr>
<tr>
<td>LDETECT_LIST_1</td>
<td>lslst1.tbl</td>
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</table>

### Note (3) compupp.epf

See file in Section B.2.

### Note (4) uslst.tbl

This symbolic link is passed as a parameter to this procedure when it is called. It points to the file containing the source list for the present declination strip. The grouped sources are indicted in Table 5.XXXX.
## Appendix B

### Abbreviations

<table>
<thead>
<tr>
<th>Abbreviation</th>
<th>Description</th>
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<tbody>
<tr>
<td>AGN</td>
<td>Active Galactic Nuclei</td>
</tr>
<tr>
<td>ASCA</td>
<td>Advanced Satellite for Cosmology and Astrophysics</td>
</tr>
<tr>
<td>AXAF</td>
<td>Advanced X-ray Astronomy Facility</td>
</tr>
<tr>
<td>BH</td>
<td>Black Hole</td>
</tr>
<tr>
<td>CDS</td>
<td>Strasbourg Astronomical Data Centre</td>
</tr>
<tr>
<td>EXSAS</td>
<td>Extended Scientific Analysis System</td>
</tr>
<tr>
<td>FOV</td>
<td>Field of View</td>
</tr>
<tr>
<td>HR</td>
<td>Hardness Ratio</td>
</tr>
<tr>
<td>HRI</td>
<td>High Resolution Imager</td>
</tr>
<tr>
<td>HST</td>
<td>Hubble Space Telescope</td>
</tr>
<tr>
<td>IPC</td>
<td>Imaging Proportional Counter</td>
</tr>
<tr>
<td>LINER</td>
<td>Low-Ionization Nuclear Emission-line Region</td>
</tr>
<tr>
<td>MC</td>
<td>Magellanic Cloud</td>
</tr>
<tr>
<td>MDO</td>
<td>Massive Dark Object</td>
</tr>
<tr>
<td>ML</td>
<td>Maximum Likelihood</td>
</tr>
<tr>
<td>MPC</td>
<td>Monitor Proportional Counter</td>
</tr>
<tr>
<td>MPE</td>
<td>Max Planck Institute for Extraterrestrial Physics</td>
</tr>
<tr>
<td>NED</td>
<td>NASA/IPAC Extragalactic Database</td>
</tr>
<tr>
<td>PET</td>
<td>Photon Event Table</td>
</tr>
<tr>
<td>PSF</td>
<td>Point Spread Function</td>
</tr>
<tr>
<td>PSPC</td>
<td>Position Sensitive Proportional Counter</td>
</tr>
<tr>
<td>ROSAT</td>
<td>Röntgen Satellite</td>
</tr>
<tr>
<td>SASS</td>
<td>Standard Analysis Software System</td>
</tr>
<tr>
<td>SIMBAD</td>
<td>Set of Identifications, Measurements and Bibliography for Astronomical Data</td>
</tr>
</tbody>
</table>
APPENDIX B. ABBREVIATIONS

SNR  Supernova Remnant
XMM  X-ray Multi-Mirror
XRT  X-ray Telescope
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