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New Methods of Studying Young Stellar Objects Using Radio Interferometry

A dissertation submitted to the University of Dublin for the degree of

Doctor of Philosophy

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Summary

While most studies of star formation are based on infrared or optical observations, radio astronomy has also proven very useful in the study of this process. This is seen for example in the detection of free-free emission from the jets of young stellar objects (YSOs), which has allowed us to estimate parameters such as the ionization fraction or the mass-loss-rate of these jets, or in the detection of synchrotron emission, indicating that electrons are accelerated to relativistic speeds in shocks of YSO jets. Thanks to the upgrading of radio interferometers such as the Karl G. Jansky Very Large Array (VLA) or the Enhanced Multi Element Remotely Linked Interferometer Network (e-MERLIN), and the resulting huge increase in sensitivity, there are new opportunities to apply radio astronomy to the study of star formation. As well as this, new, sensitive low-frequency radio interferometers such as the Low-Frequency Array (LOFAR) mean that there is now the possibility to study YSOs at much lower frequencies than previously possible.

In my first work in Chapter 2, I report on the detection of synchrotron emission in two emission knots in the jet of the low-mass YSO DG Tau A at 152 MHz using LOFAR. In one of the knots, a low-frequency turnover in its spectrum is clearly seen compared to higher frequencies. This is the first time such a turnover has been seen in nonthermal emission in a YSO jet. I consider several possible mechanisms for the turnover and fit models for each of these to the spectrum. Based on the physical parameters predicted by each model, the Razin effect appears to be the most likely explanation. From the Razin effect fit, I obtain an estimate for the magnetic field strength within the emission knot of $\sim 20 \mu\text{G}$. If the Razin effect is responsible, this

is the first time the magnetic field strength along a YSO jet has been measured based on a low-frequency turnover in nonthermal emission.

Then in Chapter 3, I report the first detection of coherent emission at low frequencies from T Tauri stars. Using LOFAR, several bright radio bursts are detected at 150 MHz from two weak-line T Tauri stars: KPNO-Tau 14 and LkCa 4. All of the bursts have high brightness temperatures ($10^{13} - 10^{14}$ K) and high circular polarisation fractions (60 - 90 %), indicating that they must be due to a coherent emission mechanism. This could be either plasma emission or electron-cyclotron maser (ECM) emission. Due to the exceptionally high brightness temperatures seen in at least one of the bursts ($\geq 10^{14}$ K), as well as the high circular polarisation levels, ECM is favoured as the most likely emission mechanism. Assuming this is the case, the required magnetic field in the emission regions would be 40 - 70 G. I determine that the most likely method of generating ECM emission is plasma co-rotation breakdown in the stellar magnetosphere. There remains the possibility, however, that it could be due to an interaction with an orbiting exoplanet.

Finally in Chapter 4, I obtain high-resolution radio images of two binary YSOs: L1551 IRS 5 and L1551 NE, using e-MERLIN, the VLA, and the Atacama Large Millimeter Array (ALMA) covering a wide range of frequencies from 5 - 336 GHz. By comparing these observations to observations at a previous epoch, it is shown that there is high degree of variability in the free-free emission from the jets of these sources. In particular, the northern component of L1551 IRS 5 shows a remarkable decline in flux density of a factor of ~ 5 . By fitting the spectra of the sources, the ionized mass loss rates of the jets are derived and it is shown that there is significant variability on timescales of several years. Thanks to the high resolution achieved at cm-wavelengths, it is also possible to study the launching and collimation zone in more detail. Using radiative transfer modelling, a model image is obtained for the jet of the southern component of L1551 IRS 5. By comparing this with the observed image, it was possible to estimate the collimation distance of the jet. This suggested the jet is collimated on scales of $\lesssim 12$ au, similar to previous upper limits.

“There is a theory which states that if ever anyone discovers exactly what the Universe is for and why it is here, it will instantly disappear and be replaced by something even more bizarre and inexplicable.

There is another theory which states that this has already happened.”

— Douglas Adams

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List of Publications

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List of Abbreviations

AGN	Active Galactic Nuclei
ALMA	Atacama Large Millimeter/sub-millimeter Array
BJD	Barycentric Julian Date
CTTS	Classical T Tauri Star
DDE	Direction-Dependent Effect
ECM	Electron-Cyclotron Maser Emission
e-MERLIN	Enhanced Multi Element Remotely Linked Interferometer Network
FFA	Free-Free Absorption
HBA	High-Band Array
HH	Herbig-Haro Object
HST	Hubble Space Telescope
GMC	Giant Molecular Cloud
GMRT	Giant Metrewave Radio Telescope
IR	Infrared
ISM	Interstellar Medium
LBA	Low-Band Array
LOFAR	Low-Frequency Array
LoTSS	LOFAR Two-Metre Sky Survey
MWA	Murchison Widefield Array
ngVLA	Next Generation Very Large Array
NVSS	NRAO VLA Sky Survey
RM	Rotation Measure

RMSF	RM Spread Function
SED	Spectral Energy Distribution
SKA	Square Kilometre Array
SSA	Synchrotron Self-Absorption
TGSS	TIFR GMRT Sky Survey
TMC	Taurus Molecular Cloud
TTS	T Tauri Star
VLA	Karl G. Jansky Very Large Array
WTTS	Weak-lined T Tauri Star
YSO	Young Stellar Object

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Chapter 1

Introduction

1.1 Motivation

One of the most important questions in astrophysics is how do stars form. The study of star formation ties into many other areas of astrophysics such as the study of the interstellar medium, galaxy structure and evolution, stellar evolution and planet formation. Star formation has primarily been studied using infrared or optical observations. However, radio astronomy has also helped us to learn more about this process. For example, a well-known and quite spectacular signpost of star formation is the phenomenon of jets and outflows from young stellar objects (YSOs). These often exhibit radio emission in the form of thermal bremsstrahlung emission thought to be from ionized plasma close to the base of the jet (Anglada et al. 2018), allowing us to estimate parameters such as the ionization fraction or the mass-loss-rate of the jet. Additionally, synchrotron emission is seen in shocks along the jet axis, indicating that these shocks are capable of accelerating electrons in the jet to relativistic speeds. Based on the polarisation of this emission we can estimate the magnetic field strength in the jet (Carrasco-González et al. 2010). As well as studies of YSO jets, in this work I show for the first time that through the detection of coherent radio bursts, the magnetic activity of YSOs can be studied and the presence of exoplanets can potentially be inferred.

In recent years, radio interferometers such as the Karl G. Jansky Very Large Array (VLA) in the USA or the Enhanced Multi Element Remotely Linked Interferometer

Network (e-MERLIN) in the UK have been upgraded to have larger bandwidths, giving a huge increase in sensitivity. Thanks to this, there is now a range of new opportunities to apply radio astronomy to the study of star formation. As well as this, new, sensitive interferometers at low frequencies such the Low-Frequency Array (LOFAR) give us the ability to study YSOs at much lower frequencies than previous radio studies which have primarily been at GHz-frequencies. Thanks to all of these developments, there has never been a better time to perform radio studies of YSOs.

Before describing the work I did in detail, I will begin by discussing the basics of the process of star formation, followed by describing the various types of radio emission observed in YSOs along with the basics of radio interferometry itself.

1.2 Star Formation

The process of star formation mostly occurs within objects known as giant molecular clouds (GMCs). These are large, dense regions in the interstellar medium (ISM) containing molecular gas. They have masses ranging from $10^2 M_\odot$ to $10^7 M_\odot$. Within the GMCs, denser regions of gas known as *clumps* are found, typically with a mass of $M \sim 10^3 - 10^4 M_\odot$. These are where groups of stars or *star clusters* are formed. Within the clumps are much denser regions known as *cores* with masses on the order of $\sim 1 M_\odot$. If a core is above a critical mass known as the *Jeans mass* M_J , the thermal energy within the core is unable to support the core against its own gravity. Therefore, it will undergo gravitational collapse and begin to form a protostar (Shu et al. 1987). The Jeans mass is given by:

$$M_J = \left(\frac{5}{2} \frac{k_B T}{m_u \mu G} \right)^{3/2} \left(\frac{4}{3} \pi \rho \right)^{-1/2} \quad (1.1)$$

where k_B is the Boltzmann constant, T is the temperature of the cloud, m_u is the atomic mass unit, μ is the molecular weight of the gas, G is the gravitational constant and ρ is the density of the gas.

Generally, star formation is divided into 3 mass regimes: low-mass star formation

covers the formation of stars with a mass M_* of $0.1 M_\odot \leq M_* \leq 2 M_\odot$, intermediate-mass star formation covers $2 M_\odot \leq M_* \leq 8 M_\odot$, and high-mass star formation covers $M_* \geq 8 M_\odot$. This work will focus on low-mass star formation.

Young stellar objects (YSOs) are stars in the process of formation. They accrete material from a surrounding circumstellar disk and envelope before eventually dissipating the disk and envelope and contracting under gravity until they become a main sequence star. They are conventionally classified based on their spectral energy distribution (SED) in the infrared at $\lambda > 2 \mu m$, using the spectral index $\alpha_{\text{IR}} \equiv d \log \lambda F_\lambda / d \log \lambda$ where F_λ is the flux density at wavelength λ . Originally, 3 classes were used, Class I, II, and III (Lada 1987). Later a fourth class, the “Flat-SED” class, was added to describe YSOs on the border between Class I and Class II (Greene et al. 1994). These are summarised as follows:

- **Class I:** $0.3 \leq \alpha$
- **Flat-SED:** $-0.3 \leq \alpha < 0.3$
- **Class II:** $-1.6 \leq \alpha < -0.3$
- **Class III:** $\alpha < -1.6$

Another class, known as Class 0, was also added to describe YSOs which could not be detected in infrared but had been detected through other means (Andre et al. 1993). These classes roughly correspond to different stages of evolution of the star. However, it is not perfect as effects such as inclination and asymmetric geometry can affect which class is observed (Robitaille et al. 2006). Typical SEDs for each of these classes can be seen in Figure 1.1.

As the molecular cloud core collapses, a protostar begins to form from material which has stopped collapsing and begun to obey quasi-hydrostatic equations. Material continues to fall on to the protostar from the surrounding envelope. Due to the conservation of angular momentum, a rotationally supported accretion disk will also begin to form around the star. In addition, we begin to see the bipolar outflow from the source. At this stage, the star is a Class 0 source as it is deeply embedded within the envelope, with most of the mass still in the envelope rather than the central protostar.

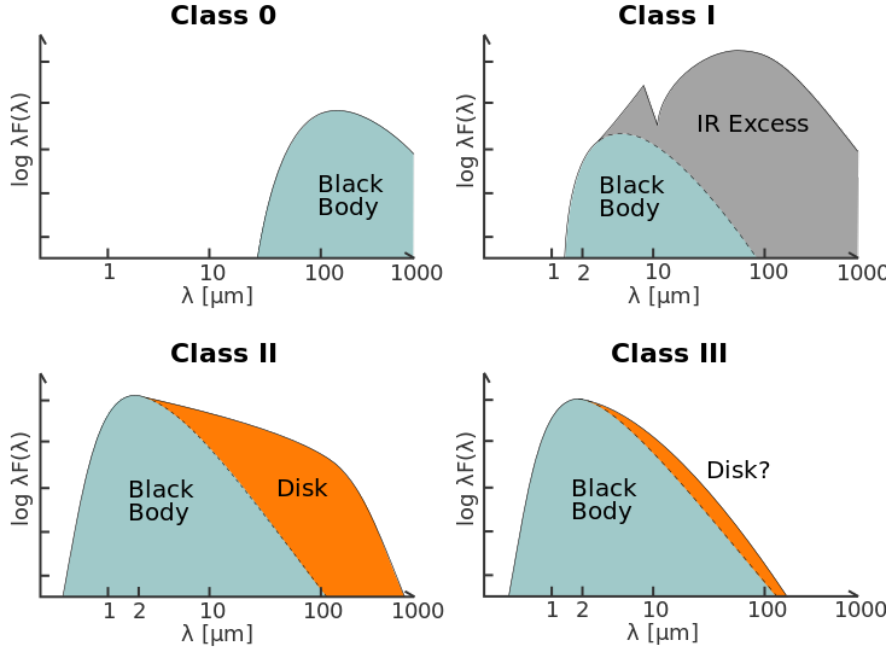


Figure 1.1: The SEDs for each of the YSO classes from (Lada 1987) and (Andre et al. 1993). Credit: Magnus Vilhelm Persson. Licensed under CC BY 4.0 (<https://creativecommons.org/licenses/by/4.0/>).

The SEDs of Class 0 objects typically look like black body with a temperature of 15–30 K. Therefore, most of its radiation is in the submillimeter or longer wavelengths. A property often used to distinguish Class 0 objects from objects at a later evolutionary phase is the bolometric temperature T_{bol} . This is defined as the temperature of a blackbody with the same flux-weighted average frequency as the observed SED, with class 0 objects having $T_{\text{bol}} < 70$ K (Dunham et al. 2014). The class 0 phase typically lasts about $\sim 1 \times 10^5$ yr (Evans et al. 2009).

As material continues to fall onto the star from the envelope, the star transitions from a Class 0 to a Class I source. At this stage, most of the mass is now in the star rather than the envelope. The SED of these objects typically peaks in the infrared as optical emission from the star is reabsorbed by the infalling envelope and then re-radiated in the infrared. For these objects, the bolometric temperature is $70 \text{ K} < T_{\text{bol}} < 650 \text{ K}$. This phase typically lasts $\sim 5 \times 10^5$ yr (Evans et al. 2009).

Eventually, as the infalling envelope begins to clear as the material is either accreted or dispersed by the outflow, the star becomes optically visible, at least if viewed from above the disk. Now it has become a Class II source. The SED of these objects

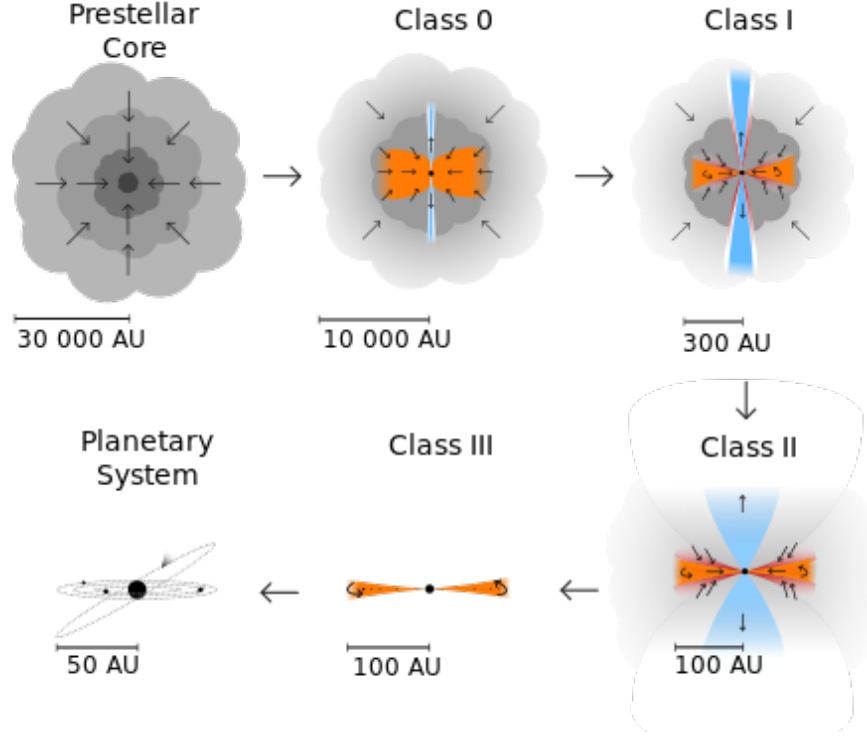


Figure 1.2: The different stages of protostellar evolution . Credit: Magnus Vilhelm Persson. Licensed under CC BY 4.0 (<https://creativecommons.org/licenses/by/4.0/>).

typically resembles a black body spectrum, but with a large infrared excess present due to the disk. The bolometric temperatures are $650 \text{ K} < T_{\text{bol}} < 2800 \text{ K}$. Accretion continues to occur onto the star from the accretion disk. However, the class observed can depend on the viewing angle. If viewed edge-on to the disk, it may still appear as a Class I source as it would be highly obscured by the disk. In the low-mass regime, Class II sources are also known as *classical T Tauri stars* (CTTSs), named after the prototypical example T Tauri.

As the star continues to evolve, it becomes a Class III source, also known as a *weak-lined T Tauri star* (WTTS) in the low-mass regime. These sources show very little or no infrared excess, with their SEDs generally resembling a black body at a single temperature, with typical bolometric temperatures of $T_{\text{bol}} > 2800 \text{ K}$. This indicates that they no longer have significant disks present. These have likely been dispersed as a result of viscous evolution and photoevaporation from the central star. As a result, these sources have finished accreting.

Once accretion is completed, the star initially produces radiation through gravita-

tional contraction. Eventually, when the pressure in the stellar core is high enough to sustain nuclear fusion reactions, the star becomes a zero-age main sequence star and the process of star formation is complete. An illustration of our current understanding of the process of protostellar evolution can be seen in Figure 1.2.

1.2.1 Jets and outflows

Astrophysical jets are a phenomenon associated with a wide variety of accreting astrophysical systems, such as YSOs, neutron stars and active galactic nuclei (AGNs). In YSOs, they are seen across the whole stellar mass spectrum, from brown dwarfs (Wheeler et al. 2005) to massive stars (Carrasco-González et al. 2010). They consist of highly collimated outflows of material ejected from the system along the axis of rotation and are believed to be launched due to a dynamic interaction between the accretion disk and the stellar and/or disk magnetic field (Frank et al. 2014). They play an important role in the evolution of stars as they carry away much of the accreting material and so help determine the final masses of stars. Typically, the ratio of jet mass flux to the accretion rate is $\approx 10\%$ (Frank et al. 2014). Additionally, they contribute to the loss of angular momentum from the star. This has been shown by the detection of rotation in the jet through the measurement of asymmetries in the radial velocity across the jet. For the outflows from the YSOs DG Tau and RW Aur, the angular momentum removed was measured to be between 60% and 100% of the angular momentum in the disk required to be removed in order for the star to accrete at the observed rate (Bacciotti et al. 2002; Woitas et al. 2005). The removal of angular momentum is essential to prevent excessive spin-up in the star as it accretes and contracts. Jets may also disperse and re-eject material in the infalling envelope, thereby reducing the final mass of the star. Through the injection of momentum and energy into the surrounding molecular cloud, jets can also play a role in feedback on the cloud. By disrupting the cloud and driving turbulence, jet feedback helps to support the cloud against collapse, and can therefore affect the percentage of the mass of the cloud converted into stars i.e. the star-formation-efficiency.

The first signs of these outflows were discovered by Herbig (1951) and Haro & Guillermo (1952), who discovered small emission-line nebulae with unusual spectra within dark clouds. Originally, these so-called Herbig-Haro objects were thought to be regions of star formation themselves. However, later on it was realized that they were in fact supersonic shocks within highly collimated jets from YSOs (Dopita et al. 1982; Mundt & Fried 1983). Visual and IR emission lines, such as the $H\alpha$ line or forbidden transitions such as [SII], [OI] and [NII], are the most commonly used tracers for jets (Frank et al. 2014). Jets can be very large in extent. Wide-field surveys in optical and IR have suggested that jets as large as several parsecs in extent are a common phenomenon. In one case, the object HH 111 shows an angular size of almost 1° on the sky at a distance of 460 pc, corresponding to a total length of 7.7 pc (Reipurth et al. 1997).

In addition to the collimated jets, there are also the wide-angle “molecular outflows”. These consist of ambient molecular gas in the molecular cloud in the vicinity of the star which has been swept up by the jet. Originally detected near the source L1551 IRS 5 (Snell et al. 1980), it was later realized to be a common feature around YSOs, with hundreds of such outflows known today (Wu et al. 2004). These outflows are more massive and less collimated than the jet and are usually traced using CO rotational lines (Frank et al. 2014).

1.2.2 Jet launching mechanism

The precise mechanism responsible for launching the jets is still a matter of great debate. Despite this, it is generally accepted that it involves magnetic coupling of accreting material with the stellar and/or disk magnetic field and magnetocentrifugal acceleration. Unfortunately, as acceleration is believed to occur within a few au of the star, resolving the launching region is difficult as the projected extent on the sky is only a few tens of milliarcseconds. However, larger scale properties of the jet, such as the ejection-to-accretion ratio, collimation and the velocity structure of the jet can give clues about the mechanism responsible (Frank et al. 2014). The exact launching

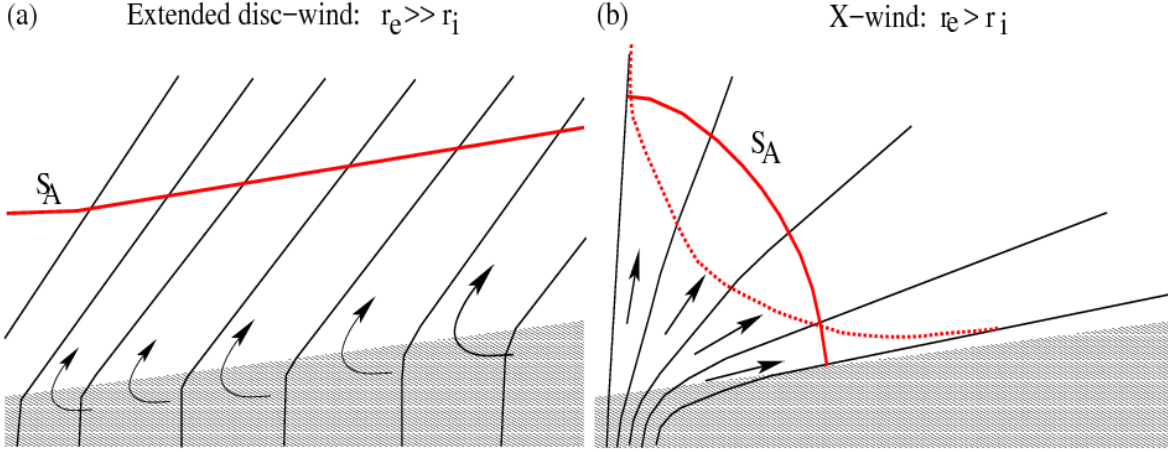


Figure 1.3: The magnetic field configurations for the jet launching mechanisms. r_i is the inner radius of the disk and r_e is the external radius, which corresponds to where the jet is launched. The Alfvén surface is indicated by S_A . Credit: Ferreira et al. (2006), reproduced with permission © ESO.

distance from the star depends on the model used.

The mechanisms proposed can be classified on the basis of where in the disk the jet originates from. Currently, there are two main competing models: the *disk wind model* and *X-wind model*. In the disk wind model, the jet is launched over a large range of radii in the disk, so that the external radius r_e where the jet is launched from is much larger than the inner radius r_i of the disk. Originally, this model was proposed as an origin for AGN jets (Blandford & Payne 1982), and was quickly adopted for YSO jets (Pudritz & Norman 1983a). A magnetic field threading the disk is required. This can be either a primordial magnetic field inherited from the parent molecular cloud, or it can be generated within the disk itself. Considering a frame of reference co-rotating with the disk at a radius R , material within the disk can become attached to the poloidal field lines. These field lines act as rigid wires for the material. If the field lines are inclined by less than a critical degree of $\theta \leq 60^\circ$ with respect to the plane of the disk, then the effective potential along the field lines *decreases* with distance. As a result, the material on the field lines will undergo what is known as magnetocentrifugal acceleration. As the material is forced to rotate at constant angular velocity by the field lines, its angular momentum increases with distance. This causes angular momentum to be extracted from the disk.

In the X-wind model (Shu et al. 1994), the jet is launched from close to the star, so r_e is only slightly bigger than r_i . The magnetic field involved is generated by the star. The poloidal component of the magnetic field is given by a dipole aligned with the rotation axis of the star and is stationary in a frame that rotates with the star. The magnetic field truncates the accretion disk out to some radius r_t . r_t is generally somewhat smaller than the co-rotation radius r_x , where the angular velocity Ω_x of the disk is equal to the angular velocity Ω_* of the star's rotation at its surface. Just outside of r_t the dipole field lines are pinched inwards by the accretion flow towards the star. This causes the magnetic field lines outside r_x to be bent outwards. The magnetic field enforces corotation of the disk at the angular velocity Ω_x in the region between r_t and slightly beyond r_x . As a result of this, in a region known as the T-region between r_t and r_x , the gas in the disk will be orbiting at sub-Keplerian velocities. The gas then attaches to the field lines which bow inwards toward the star and fall onto the star. This is known as the funnel flow. Slightly outside r_x in a region known as the X region, the gas is rotating at super-Keplerian velocities. The gas will then attach to the field lines which bow outwards and move outwards along them. The particles are accelerated along the field lines by the magnetocentrifugal force similar to the disk wind model.

In both models, the material will eventually be accelerated to the Alfvén speed, once it reaches a location known as the Alfvén surface. Above this speed, the magnetic field lines can no longer control the flow and no longer act as rigid wires. Therefore, magnetocentrifugal acceleration ceases and the angular momentum in the flow remains constant beyond that point.

Once the jet has been accelerated, observations suggest that collimation occurs within approximately 100 au of the source (Frank et al. 2014). Initially, it was suggested that a pressure gradient in the surrounding infalling envelope could act as a collimation mechanism. However, observations show that collimation occurs on similar scales in Class 0 and Class I objects as it does in Class II objects (Cabrit et al. 2007; Davis et al. 2011), suggesting that the infalling envelope does not play an important role

in collimation. Instead, the most likely mechanism seems to be collimation by the magnetic field responsible for launching the jet. Beyond the Alfvén surface, although the poloidal component of the field is unimportant, the toroidal component becomes important and generates an inward force on the flow known as a “hoop stress” which collimates the jet.

1.3 Radio Emission from YSOs

YSOs are frequently observed to have emission in the radio continuum. Common sources of radio emission observed include thermal free-free emission seen from the base of the ionized jet, and thermal dust emission seen from the accretion disk. Non-thermal emission is also observed from YSOs. Usually, this is in the form of gyrosynchrotron emission from mildly relativistic electrons in the stellar magnetosphere. Synchrotron emission from highly relativistic electrons has also been observed in radio knots along the jet, at large distances from the core, moving away at high velocities (Anglada et al. 2018). In addition, coherent radio emission has been detected from YSOs in this work (see Chapter 3) and on at least two previous occasions (Slee & Higgins 1971; Smith et al. 2003), in the form of bursts on time-scales of minutes to hours.

Before describing the various radio emission mechanisms in more detail, several useful quantities will be defined. The *specific intensity* or *brightness* I_ν of radiation is defined as the infinitesimal power dP intercepted by an infinitesimal area $d\sigma$ from an infinitesimal area on the sky $d\Omega$ in an infinitesimal bandwidth $d\nu$:

$$I_\nu = \frac{dP}{\cos\theta d\Omega d\sigma d\nu} \quad (1.2)$$

where θ is the angle between the normal to $d\sigma$ and the direction to $d\Omega$.

For a black body in thermal equilibrium, the brightness distribution is given by the Planck function which depends only on the temperature T of the black body:

$$B_\nu(T) = \frac{2h\nu^3}{c^2} \frac{1}{e^{h\nu/k_B T} - 1} \quad (1.3)$$

where h is the Planck constant, c is the speed of light, and ν is the frequency of the emission. If $h\nu \ll kT$, as is generally the case at radio frequencies, this can be simplified to give what is known as the Rayleigh-Jeans law:

$$B_\nu(T) \approx \frac{2\nu^2}{c^2} k_B T \quad (1.4)$$

Therefore, if the source of radiation can be approximated as a black body, such as in a source with a large optical depth, then $I_\nu \approx B_\nu(T)$. However, this does not apply in most situations and so I_ν will differ from $B_\nu(T)$.

In radio astronomy, the brightness of sources is often measured using the brightness temperature T_B , which is defined based on the Rayleigh-Jeans law:

$$T_B = \frac{c^2}{2k_B\nu^2} I_\nu \quad (1.5)$$

In other words, it is the temperature of a black body which would generate the same measured brightness as the source.

Similarly, we can define another quantity based on the source function $J_\nu = \eta_\nu/\kappa_\nu$, where η_ν is the emissivity and κ_ν is the absorption coefficient. This quantity is known as the effective temperature T_{eff} :

$$T_{\text{eff}} = \frac{c^2}{2k_B\nu^2} J_\nu \quad (1.6)$$

In the case of a Maxwellian distribution of electrons with a temperature of T , T_{eff} will always be equal to T .

These quantities can be substituted into the equation of radiative transfer:

$$\frac{dI_\nu}{ds} = -\kappa_\nu I_\nu + \eta_\nu \quad (1.7)$$

to give the form:

$$\frac{dT_B}{d\tau_\nu} = T_B - T_{\text{eff}} \quad (1.8)$$

where τ_ν is the *optical depth* defined by $d\tau_\nu = -\kappa_\nu ds$. This can be rearranged as an integral to give:

$$T_B = \int_0^{\tau_\nu} T_{\text{eff}} \exp(-\tau_\nu) d\tau_\nu + T_B \exp(-\tau_\nu) \quad (1.9)$$

In the case of an isolated source where T_{eff} is constant, this can be simplified to:

$$T_B = T_{\text{eff}}(1 - \exp(-\tau_\nu)) \quad (1.10)$$

This relation has an important consequence, as it shows that for incoherent radiation, T_B will always be less than T_{eff} . As a result, very high brightness temperatures generally imply a coherent radiation mechanism.

1.3.1 Free-free Emission

Free-free emission, also known as *bremsstrahlung* emission is a type of radiation emitted by charged particles decelerating in the electromagnetic field of another charged particle and losing energy, typically produced by free electrons in a plasma being scattered by ions, and is commonly observed in YSOs, particularly in younger Class 0 and Class I objects. The emission coefficient is given by (Rybicki & Lightman 1979):

$$j_\nu \propto T_e^{-1/2} n_e^2 e^{-h\nu/k_B T} \quad (1.11)$$

where T_e is the electron temperature and n_e is the electron density. There is also an inverse absorption process present known as free-free absorption (FFA), where a free electron can gain energy from an ion by absorbing a photon. The absorption coefficient for FFA at low frequencies is given by (Rybicki & Lightman 1979):

$$\alpha_\nu^{\text{FFA}} \propto T_e^{-3/2} n_e^2 \nu^{-2} \quad (1.12)$$

For a plasma of uniform density, the resulting radio spectrum will consist of two components, an optically thick and an optically thin component (Scaife 2013). At low frequencies, the source is optically thick ($\tau_\nu \gg 1$) and the free-free emission is given approximately by a power law of the form $S_\nu \propto \nu^2$. Above a threshold frequency ν_m , defined as the frequency where $\tau_\nu = 1$, the source will become optically thin and the emission is given by $S_\nu \propto \nu^{-0.1}$. However, in practice, sources are not uniform in density and the emission observed will come from a mixture of optically thick and optically thin regions, resulting in a shallower spectrum ($0 \leq \alpha \leq 2$).

In YSOs, this emission is found to be elongated in the direction of the outflow of the source (Anglada et al. 2018). Based on the form of their spectra and their

morphology, these radio sources are thought to be a result of free-free emission from a thermally ionized jet close to the source (Reynolds 1986) and are commonly referred to as “thermal radio jets”. Generally, this emission is detected with spectral indices of $\sim -0.1 \leq \alpha \leq 1$, with a median value of 0.45, where the spectral index α is defined by $S_\nu \propto \nu^\alpha$, where S_ν is the flux density and ν is the frequency (Anglada et al. 2018). Close to the source, the jet will be optically thick, while further out the jet will become optically thin. Therefore, depending on the frequency and the inclination of the jet, a combination of optically thick and optically thin emission will be observed.

Reynolds (1986) modelled the emission observed from an ionized jet and determined that it should depend on the geometry of the jet. This model assumes that the ionized jet begins at an injection radius r_0 , has an initial half-width w_0 with a circular cross-section, and the half-width of the jet has a power-law dependency given by:

$$w = w_0 \left(\frac{r}{r_0} \right)^\epsilon \quad (1.13)$$

For an isothermal jet with a constant velocity and ionization fraction, the spectral index will then be given by:

$$\alpha = 1.3 - \frac{0.7}{\epsilon} \quad (1.14)$$

For a typical, collimated, ionized jet, $\alpha \approx 0 - 1$. Above a threshold frequency ν_m however, the entire jet will become optically thin and the spectral index then becomes $\alpha = -0.1$. Thermal radio emission from YSOs is generally dominated by free-free emission at frequencies $\lesssim 20$ GHz as shown in Figure 1.4.

Given that YSOs are often heavily obscured at shorter wavelengths such as optical and IR, observations at radio wavelengths have the advantage that they do not suffer from extinction and therefore allow ejection from the system close to the central object to be traced. Using this, the direction and collimation of the material recently ejected can be observed. It also allows physical properties of the jet, such as the ionized mass loss rate $\dot{M}_{\text{ion}}$, to be measured. Typically, when comparing $\dot{M}_{\text{ion}}$ with the total mass loss rate $\dot{M}$ obtained from the large scale molecular outflow or from atomic emission

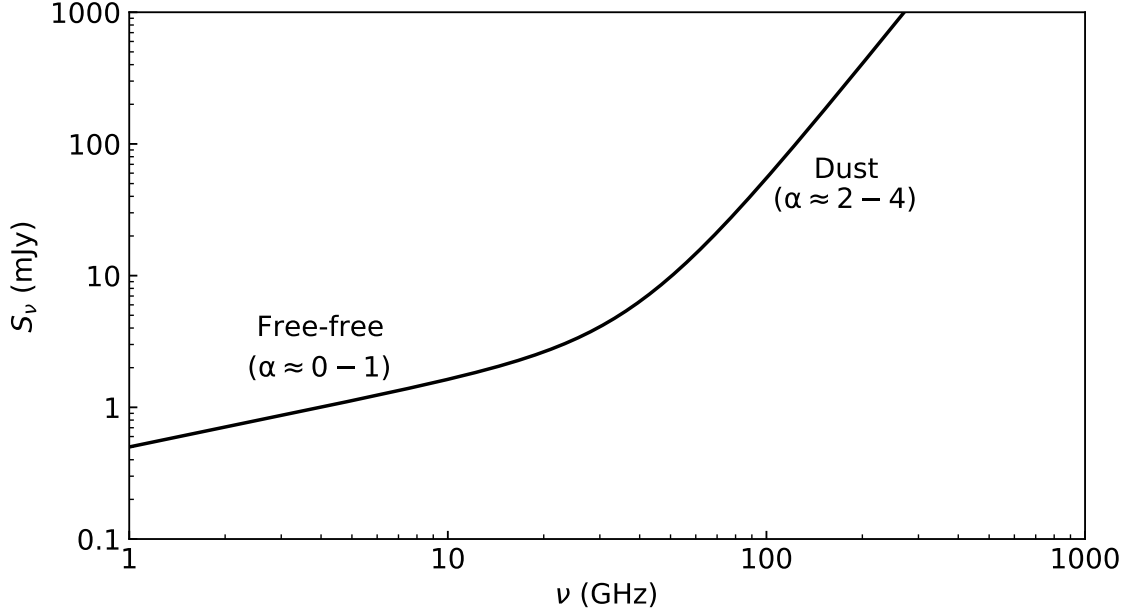


Figure 1.4: A typical SED for a YSO source emitting free-free emission and thermal dust emission

lines, $\dot{M}_{\text{ion}}$ is an order of magnitude smaller than $\dot{M}$ (Anglada et al. 2018). This implies that the jets are only partially ionized, with an ionization fraction of $\sim 1 - 10\%$.

1.3.2 Thermal Dust Emission

Radio emission in YSOs can also result from dust emission in the disks surrounding the stars. Dust emission is heavily dominated by modified blackbody emission (Scaife 2013). In the Rayleigh-Jeans limit and optically thin limit, this can be characterised as:

$$S_\nu \propto \nu^{\beta/(1+\Delta)} B_\nu(T_d) \quad (1.15)$$

where β is the opacity index of the dust particles, Δ is the ratio of optically thick to optically thin emission, and T_d is the dust temperature. The spectral index can then be related to the opacity index as $\alpha = 2 + \beta/(1 + \Delta)$. Typically, dust emission is characterised by spectral indices of $\alpha \approx 2 - 4$. The value of β and therefore of α depends on the dust composition and grain size distribution. Smaller grains have larger opacity indices leading to steeper spectra, while larger grains have smaller opacity indices leading to flatter spectra. In radio emission from YSOs at frequencies $\gtrsim 20$

GHz, dust emission tends to dominate over bremsstrahlung emission (see Figure 1.4).

1.3.3 Non-Thermal Emission

Non-thermal emission in YSOs is generally attributed to gyrosynchrotron emission and synchrotron emission. Both of these are caused by electrons in a plasma spiralling in a magnetic field, with gyrosynchrotron emission being due to mildly relativistic electrons ($\gamma \sim 2 - 3$) and synchrotron being due to ultra-relativistic electrons ($\gamma \gg 1$).

Gyrosynchrotron emission from YSOs is generally compact and centred on the star and is thought to be produced by coronal-type activity. It can show high levels of variability on timescales of weeks to months, with some extreme sources even showing variability on timescales as short as 30 minutes (Forbrich et al. 2017). Gyrosynchrotron sources usually have spectral indices of $\alpha < -0.1$ and often show circular polarization with polarization fractions of $\lesssim 30\%$ (Dzib et al. 2013; Dzib et al. 2015). While centimetre radio emission from YSOs tends to be dominated by thermal bremsstrahlung emission from radio jets in Class 0 and Class I, gyrosynchrotron emission usually dominates in Class III objects, with Class II objects showing a mixture of both emission types.

Synchrotron emission in YSOs is usually found along the jet at large distances from the star. Based on this, it is believed to be tracing strong shocks of the jet against the surrounding molecular cloud or internal working surfaces in the jet resulting from variable ejection events (Purser et al. 2018). Particles in the jet are believed to be accelerated within the shocks to relativistic velocities through the process of diffusive shock acceleration (DSA) (Bell 1978), where the particles gain energy by diffusing back and forth across the shock front.

In synchrotron emission, the electron distribution is usually in the form of a non-Maxwellian power-law distribution:

$$n(E) = KE^{-p} \quad (1.16)$$

where K is a constant and E is the energy of the electrons. The emission is distributed over a broad continuum at high harmonics of the electron-cyclotron frequency ν_B and is strongly directed in the direction of the electron's instantaneous motion. For an isotropic electron distribution, this means that the emission peaks perpendicular to the magnetic field lines. The emissivity coefficient is given by (Rybicki & Lightman 1979):

$$j_\nu \propto \kappa B^{(p+1)/2} \nu^{-(p-1)/2} \quad (1.17)$$

where B is the magnetic field strength.

Similar to bremsstrahlung, it also has an inverse absorption process known as synchrotron self-absorption (SSA), where photons interact with electrons in the magnetic field and are absorbed. The absorption coefficient for SSA is given by:

$$\alpha_\nu^{\text{SSA}} \propto \kappa B^{(p+2)/2} \nu^{-(p+4)/2} \quad (1.18)$$

For optically thick emission, the spectral index is 2.5, while for optically thin emission, it will depend on the power law index p of the electron distribution:

$$\alpha = -(p - 1)/2 \quad (1.19)$$

Currently the best studied cases of synchrotron emission are the intermediate-mass YSO Serpens and the high-mass YSO HH80-81. HH80-81 was the first and so far only YSO to have linearly polarized emission detected in its jet (Carrasco-González et al. 2010), as shown in Figure 1.5, confirming the presence of synchrotron emission and therefore of relativistic particles. This seems to support DSA as the origin of the emission. This also allowed the structure and strength of the magnetic field in the jet to be measured. The direction of polarization was found to be perpendicular to the jet with the degree of polarisation increasing with distance from the jet axis, interpreted as a helical magnetic field structure in the jet. The magnetic field strength in the jet was estimated using the classical equipartition method (Beck & Krause 2005a), which assumes an equipartition of energy between the magnetic field and the relativistic

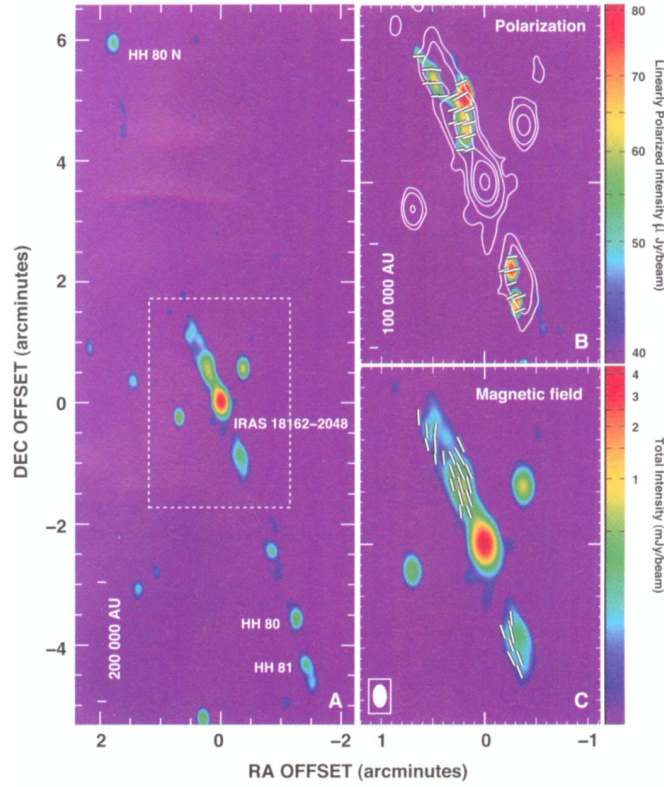


Figure 1.5: HH80-81 at 6 cm wavelength as observed by Carrasco-González et al. (2010). (A) The total continuum flux density image showing the whole extension of the jet. The brightest knot labelled IRAS 18162-2048 is thermal bremsstrahlung emission associated with the central protostar. Emission knots due to synchrotron emission in shocks of the jet, labelled HH 80, HH 81, and HH 80 N are seen further out along the jet. (B) A close-up of the inner region of the jet showing the linear polarized emission shown in colourscale. The total emission from (A) is shown as contours. The direction of the linear polarization is shown as white bars. (C) The magnetic field direction is shown as white bars superimposed on top of the total emission colourscale image. Credit: Carrasco-González et al. (2010). Reprinted with permission from AAAS.

particles, and was found to be $B \approx 200 \mu\text{G}$, at a distance of 0.5 pc from the star.

Both HH80-81 and Serpens were observed recently with the VLA in the 1-10 GHz range and with high angular resolution (Rodríguez-Kamenetzky et al. 2016, 2017). These observations showed non-thermal emission originating in shocks at termination points of the jet. In the case of HH80-81, it was also found that there were regions of alternating positive and negative spectral indices, suggesting that internal shocks within the jet are also capable of accelerating particles to relativistic velocities.

1.3.4 Coherent Emission Mechanisms

Two types of coherent emission mechanisms, in which particles emit in phase, are commonly seen in stellar radio emission: plasma emission and electron-cyclotron maser (ECM) emission.

Electron-Cyclotron Maser Emission

In ECM emission, radiation is amplified at harmonics of the electron-cyclotron frequency ν_B :

$$\nu_B = \frac{\Omega_e}{2\pi} = \frac{eB}{2\pi m_e c} \quad (1.20)$$

where B is the local magnetic field in the plasma. This type of emission is seen in the decametric (DAM) radio emission of Jupiter, the Auroral Kilometric Radiation (AKR) emission from Earth, and has been seen in stellar emission from various stars. The theory of this mechanism was first proposed by Twiss (1958). In order for this mechanism to operate, the following conditions need to be met. There must be a population inversion in the electron distribution compared to equilibrium, i.e. a pump for the maser. In addition, the condition $\nu_B > \nu_p$ must be met. Otherwise most of the energy will be dispersed in the form of plasma waves rather than electromagnetic. Therefore, either a relatively strong magnetic field or a low density plasma is required.

The most common type of population inversion which is observed is what is known as a loss-cone distribution (Dulk 1985). This is produced when electrons are propagating along a magnetic field with converging foot points. This can occur for example in a planetary magnetosphere or in the magnetic field lines of an active region on a star.

To produce a loss-cone distribution, the electrons in the plasma must first be accelerated, for example as result of magnetic reconnection, such that they have a component parallel and perpendicular to the field. As the electrons propagate along the magnetic field towards the converging footpoints of the field lines, electrons with small pitch angles α (where $\alpha = \arctan(v_\perp/v_\parallel)$) below a certain critical value α_* will precipitate into the atmosphere and are lost. On the other hand, due to the conservation of magnetic moment, electrons with $\alpha > \alpha_*$ will be reflected at the foot points in what is known as

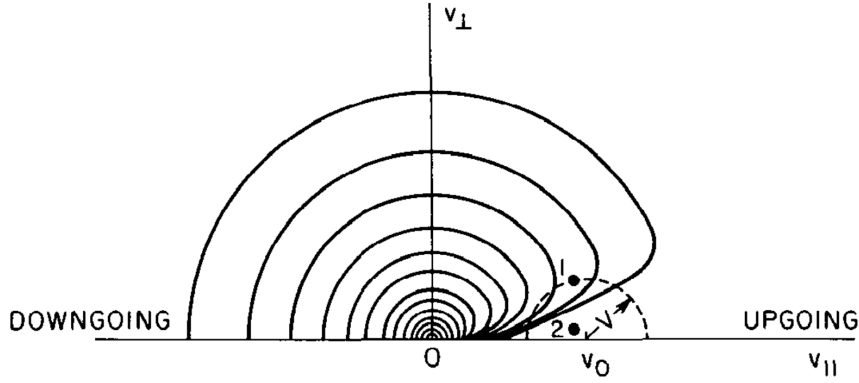


Figure 1.6: Contour map of an electron velocity distribution in $(v_{\parallel}, v_{\perp})$ space, showing an idealized loss-cone distribution. The dense semi-circular contours near the center represent the isotropic background electron distribution, while the outer ellipses deviating from a semi-circular distribution show the loss-cone. An electron at position 1, which is induced by a wave to move to position 2, would produce induced emission, while there would be few electrons at position 2 that could move to position 1 and absorb emission, leading to net amplification of the wave. Credit: (Dulk 1985). Reprinted with permission from Annual Reviews.

magnetic mirroring. This leads to the resulting loss-cone distribution. The value of α_* depends on the ratio of the magnetic field at the top of the magnetic field line and at the footpoint:

$$\alpha_* = \arcsin(B_{\text{top}}/B_{\text{foot}}) \quad (1.21)$$

The resulting electron velocity distribution can be seen in Figure 1.6. As can be seen, there is a strong deviation from a normal isotropic distribution as there is a lack of upgoing electrons with small pitch angles i.e. a large $v_{\perp}$ and a small $v_{\parallel}$. This is the source of the population inversion. Consider an electron of energy E_1 at position 1 in Figure 1.6. If this electron is induced by an electromagnetic wave to move to position 2, where $E_2 < E_1$, then this will create induced emission. On the other hand, there are few electrons at position 2 that can move to position 1 and absorb emission. Therefore this leads to net amplification of the wave.

The emission produced will be at the local electron-cyclotron frequency ν_B . While the bandwidth for a single pulse of ECM emission is small, the bandwidth of the overall emission tends to be quite large as the emission should be generated at a range of heights along the magnetic field line and thus at a range of B and ν_B values. The emission will also be very highly beamed at an angle of $\theta \approx 70^\circ$ to the magnetic field

lines and an angular range of $\approx 10^\circ$.

A major issue with ECM emission is gyroresonance absorption. In a stellar atmosphere, virtually all of the radiation produced at a frequency $\omega \approx \Omega_e$ will be absorbed at the second harmonic layer in the atmosphere where $B = B_{\text{src}}/2$. As a result, it seems unlikely that first harmonic ECM radiation can escape the atmosphere.

To resolve this problem, several solutions have been proposed. The most obvious is that radiation is produced at the second harmonic ($\omega \approx 2\Omega_e$), to avoid having to cross the second harmonic layer. However, the emission would still have to cross the third harmonic layer (where $B = 2B_{\text{src}}/3$). The optical depth would therefore still not be negligible ($\tau \approx 10$) but conditions such as an unusually low ambient density or temperature could allow a significant amount of radiation to escape.

Plasma Emission

In plasma emission, radiation is generated at the plasma frequency ν_p and its second harmonic (Dulk 1985):

$$\nu_p = \frac{\omega_p}{2\pi} = \left(\frac{n_e e^2}{\pi m_e} \right)^{1/2} \quad (1.22)$$

Here we assumed for simplicity that no magnetic field is present. This form of emission is frequently seen from the Sun in the form of Type I, II, and III bursts, for example.

The origin of plasma emission involves two stages. First, production of longitudinal (Langmuir) waves in the plasma and secondly, the conversion of these Langmuir waves into electromagnetic waves which can escape the plasma.

In order to generate these Langmuir waves, there must be nonequilibrium electron velocity distribution to provide a source of free energy, similar to how ECM emission is generated. In the case of plasma emission, this is thought to take the form of a “bump-in-the-tail” distribution, where there is an excess of electrons in velocity space at some velocity v_0 , such that there is a positive slope in the electron distribution at $v \leq v_0$. In addition, v_0 must be much larger than the background thermal velocity of the electrons (i.e. $v_0 \gg v_e = (kT/m)^{1/2}$), so that there is a negligible number of ambient electrons at v_0 . In the Sun, typically $v_0 \approx 0.1c - 0.5c$.

To generate such a distribution, electrons must be accelerated to high speeds. Moving away from the acceleration region, faster electrons will outpace the slower electrons. As a result, at a sufficiently far distance, there will be more faster electrons. This is generally thought to occur through the injection of heated plasma into a cooler ambient plasma.

Some portion of the energy of the Langmuir waves will then be converted to electromagnetic waves. Fundamental emission is generated by the scattering of Langmuir waves off of thermal ions while second harmonic emission is generated by the coalescence of two Langmuir waves travelling in opposite directions.

First harmonic plasma emission is expected to be $\approx 100\%$ polarised and to propagate in the ordinary mode (o-mode). The second harmonic however, is observed to be only weakly polarised ($\lesssim 20\%$) in either x- or o-mode. However, Vedantham (2021) found that with high magnetic field strengths, such as those seen in late-type stars, polarization fractions of up to $\approx 75\%$ could in theory be achieved under certain conditions. For both fundamental and second harmonic emission, the observed polarization fraction will likely be lower due to geometric and propagation effects. For example, the superposition of oppositely polarized emission from positive and negative line-of-sight magnetic field components will reduce the observed polarization.

While plasma emission can suffer from free-free absorption, this is generally not an issue at low frequencies as the absorption is proportional to frequency ($\kappa_\nu \propto \nu^2$). However, as a result, plasma radiation becomes increasingly rare at higher frequencies.

1.3.5 Low-Frequency Emission

Most radio observations of YSOs have been carried out at high frequencies ($\nu \geq 1$ GHz). This is largely due to the poor sensitivity of low-frequency interferometers in the past. Most emission from YSOs is thermal, whose flux density increases with frequency. As a result, the thermal emission at low frequencies is usually too weak to be detected. Thermal emission was observed at 323 MHz and 608 MHz in the YSO jets L1551 IRS5, T Tau and DG Tau A using the GMRT (Ainsworth et al. 2016). Recently, free-free

emission was also observed at 149 MHz from T Tau using LOFAR (Coughlan et al. 2017), making this the lowest frequency observation of thermal emission in a YSO jet to date. A low-frequency turnover was observed in the free-free spectrum, which allowed the degeneracy between emission measure and electron density to be broken.

For non-thermal emission, the flux density decreases with frequency. Therefore, low-frequency observations could be an excellent way of detecting such emission as it should become easier to detect at low frequencies. Despite this, few such observations have been carried out so far, with most detections of non-thermal emission being at higher frequencies. Non-thermal emission was detected from DG Tau A in the previously mentioned GMRT observations (Ainsworth et al. 2014). HH80-81 was also observed using GMRT at 325 MHz and 610 MHz (Vig et al. 2018). These observations were used to refine the estimates for the magnetic field strength in the jet, which was found to be in the range $116 - 180 \mu\text{G}$.

1.4 Radio Interferometry

A telescope has two basic properties which determine how powerful it is: its sensitivity, which is its ability to collect light, and its angular resolution, which is its ability to distinguish different sources of radiation. Both of these are improved by increasing the aperture of the telescope.

The angular resolution of a telescope is the minimum separation two point sources must have in order to be distinguished. For a single dish telescope, the angular resolution is given approximately by $\sim \lambda/D$ where λ is the wavelength of the radiation being observed and D is the diameter of the telescope. At optical wavelengths, to get a resolution of 1 arcsecond, a telescope would therefore need a diameter of 1.4 m. However, for a radio telescope observing at 1 m, in order to have the same resolution it would need a diameter of 250 km. This is a huge problem for radio astronomy.

In order to overcome this issue the technique of interferometry is used (Wilson et al. 2013), whereby two or more telescopes are combined and the interference pattern of the signals from the telescopes is obtained in order to achieve higher resolution. Using this technique, the same angular resolution can be achieved with two telescopes separated by a baseline b as with one telescope with a diameter as large as b i.e. an angular resolution of $\sim \lambda/b$, a huge improvement over a resolution of $\sim \lambda/D$.

1.4.1 Antenna Fundamentals

An antenna is a device which converts an electromagnetic wave in space into an electric current in a conductor or vice versa, depending on if it is a receiving or conducting antenna. Radio telescopes are made up of receiving antennas.

The simplest example of an antenna is a dipole antenna. This is the type of antenna used in LOFAR for example. Most radio telescopes use reflectors to focus the radiation onto a feed antenna. This is seen in the VLA or ALMA for example.

Several properties of antennas will first be defined here. For a transmitting antenna, the *power pattern* P is the angular distribution of its radiated power. Often, it is normalized to peak of the distribution, giving the *normalized power pattern*

$P_n = (1/P_{\max})P(\theta, \phi)$. Generally, most characteristics of a transmitting antenna are the same as for an equivalent receiving antenna, and vice versa. Therefore, P also describes how sensitive a receiving antenna is to radiation from different parts of the sky. For a typical antenna, P peaks in a certain part of the sky known as the *main lobe*, with much lower varying values of P seen in the remainder of the sky. Smaller peaks in P are known as *side lobes*. A typical power pattern including the main lobe and side lobes is illustrated in Figure 1.7.

An important property of a receiving antenna is the *effective area*, which describes how efficient an antenna is at converting the radiation falling on it to electrical power. It is defined as:

$$A_e = \frac{2P_\nu}{S_\nu} \quad (1.23)$$

where S_ν is the flux density of an unpolarized point source and P_ν is the output power of the antenna. The factor of 2 is due to the fact that an individual antenna can only detect one polarization and so can only detect half of the radiation from an unpolarized source. The *average collecting area* is the average of A_e accross the whole sky:

$$\langle A_e \rangle = \frac{\int_{4\pi} A_e(\theta, \phi) d\Omega}{\int_{4\pi} d\Omega} = \frac{1}{4\pi} \int_{4\pi} A_e(\theta, \phi) d\Omega \quad (1.24)$$

It can be shown that $\langle A_e \rangle$ is given by:

$$\langle A_e \rangle = \frac{\lambda^2}{4\pi} \quad (1.25)$$

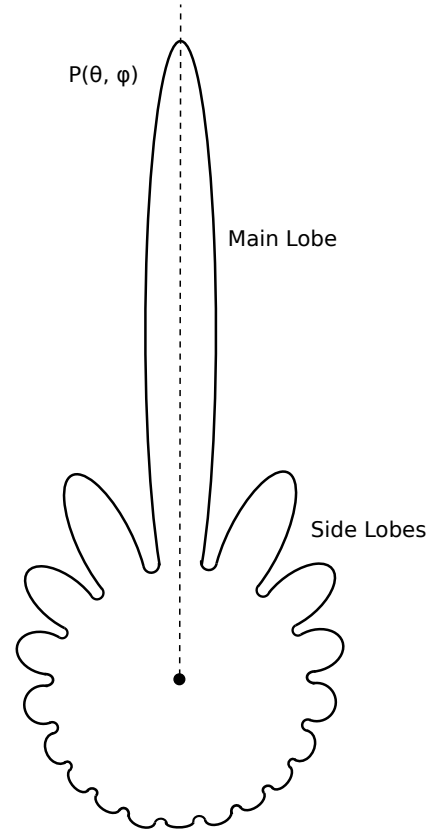


Figure 1.7: Typical power pattern for an antenna showing the main lobe and side lobes

From this, it can be seen that an isotropic antenna, such as a dipole, will become more sensitive at longer wavelengths. Therefore, dipoles are often used in radio telescopes

observing at low frequencies. However, at higher frequencies, they are generally not sensitive enough and so reflector antennas are used instead.

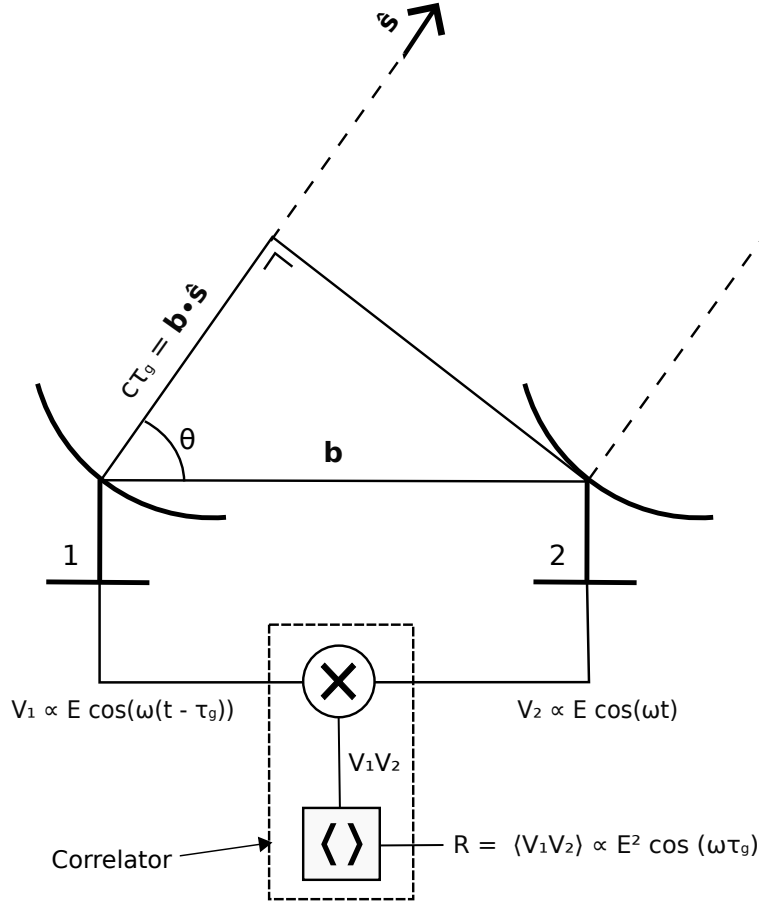


Figure 1.8: Basic elements of a two-component multiplying interferometer.

1.4.2 Two-element Radio Interferometer

Any interferometer consisting of N antennas can be treated as $N(N - 1)/2$ pairs of antennas. Therefore, it is useful to look at the case of a two-element interferometer, as seen in Figure 1.8. If the two antennas are separated by a baseline $\mathbf{b}$ and are pointed at a source with a direction given by the unit vector $\mathbf{s}$, then the signal will arrive at one antenna at a time τ_g , known as the geometric delay, later than the other antenna. This delay is given by:

$$\tau_g = \frac{\mathbf{b} \cdot \mathbf{s}}{c} = \frac{b \cos \theta}{c} \quad (1.26)$$

A plane electromagnetic wave with an amplitude E and angular frequency ω will

then induce voltages at the antennas of:

$$V_1(t) \propto E \cos(\omega(t - \tau_g)) \quad V_2(t) \propto E \cos(\omega t) \quad (1.27)$$

In an interferometer, the two signals are then correlated by passing them through a correlator, which first multiplies the signals by each other:

$$V_1(t)V_2(t) \propto E^2(\cos(2\omega t - \omega\tau_g) + \cos(\omega\tau_g)) \quad (1.28)$$

and then averages over time to remove the high frequency component $\cos(2\omega t - \omega\tau_g)$, giving a correlator output of:

$$R(t) = \langle V_1(t)V_2(t) \rangle \propto E^2 \cos(\omega\tau_g) \quad (1.29)$$

As the Earth rotates and the direction of the source changes, τ_g will vary slowly with time and so the output will oscillate quasi-sinusoidally. These sinusoidal variations are known as fringes. The fringe phase is:

$$\phi = \omega\tau_g = \frac{\omega b \cos \theta}{c} \quad (1.30)$$

The fringe phase has a dependence on θ of:

$$\frac{d\phi}{d\theta} = \frac{\omega b \sin \theta}{c} = 2\pi \left(\frac{b \sin \theta}{\lambda} \right) \quad (1.31)$$

This means that the fringe period $\Delta\phi = 2\pi$ corresponds to an angle change of $\Delta\theta = \frac{\lambda}{b \sin \theta}$. This means that the fringe phase can measure the position of a radio source very accurately if the projected baseline $b \sin \theta$ is many wavelengths long.

Also note that the factor E^2 in Equation 1.29 is proportional to the energy of the incident electromagnetic wave. Therefore the amplitude of the correlator output is proportional to the power received by the antenna.

1.4.3 Complex Visibility

When measuring a spatially extended brightness distribution $I(\sigma)$, the response of a “cosine” correlator is obtained by integrating over the source and is given by (Condon & Ransom 2016):

$$R_C = \int \mathcal{A}(\sigma) I(\sigma) \cos(2\pi \nu \mathbf{b} \cdot \sigma / c) d\Omega \quad (1.32)$$

where σ is given by $\mathbf{s} = \mathbf{s}_0 + \sigma$, and where $\mathcal{A}(\sigma) \equiv A_e(\sigma)/A_0$ is the normalized effective area in the direction σ , A_0 being the peak effective area of the antenna. $\mathbf{s}_0$ is a vector representing a position that is known as the *phase tracking center*. This is the position that is specified to be the center of the field of view. However, the cosine correlator is only sensitive to the even part of the brightness distribution, which has an even and odd part $I(\sigma) = I_E(\sigma) + I_O(\sigma)$.

To measure the odd component, a “sine” correlator is needed. To do this, a second correlator is added which has a $\pi/2$ phase delay in one of the antennas so it gives an output $R_S \propto E^2 \sin(\omega t)$. This gives a response of:

$$R_S = \int \mathcal{A}(\sigma) I(\sigma) \sin(2\pi \nu \mathbf{b} \cdot \sigma / c) d\Omega \quad (1.33)$$

The cosine and sine correlator can be combined into what is known as a complex correlator, by using Euler’s theorem to write them as a complex exponential $e^{i\phi} = \cos(\phi) + i \sin(\phi)$. The response of an interferometer with a complex correlator to an extended source is known as the *complex visibility* and is given by (Condon & Ransom 2016):

$$V \equiv R_C - iR_S = |V| e^{i\phi} = \int \mathcal{A}(\sigma) I(\sigma) e^{-2\pi i \nu \mathbf{b} \cdot \sigma / c} d\Omega \quad (1.34)$$

where $|V| = (R_C^2 + R_S^2)^{1/2}$ is known as the *visibility amplitude* and $\phi = \tan^{-1}(R_C/R_S)$ is known as the *visibility phase*. This is a complex quantity which has the same units as the flux density ($\text{W m}^{-2} \text{ Hz}^{-1}$).

The visibility is usually written in the (u, v, w) co-ordinate system, where the baseline $\mathbf{b}$ is measured using the vectors (u, v, w) which point in the directions east, north

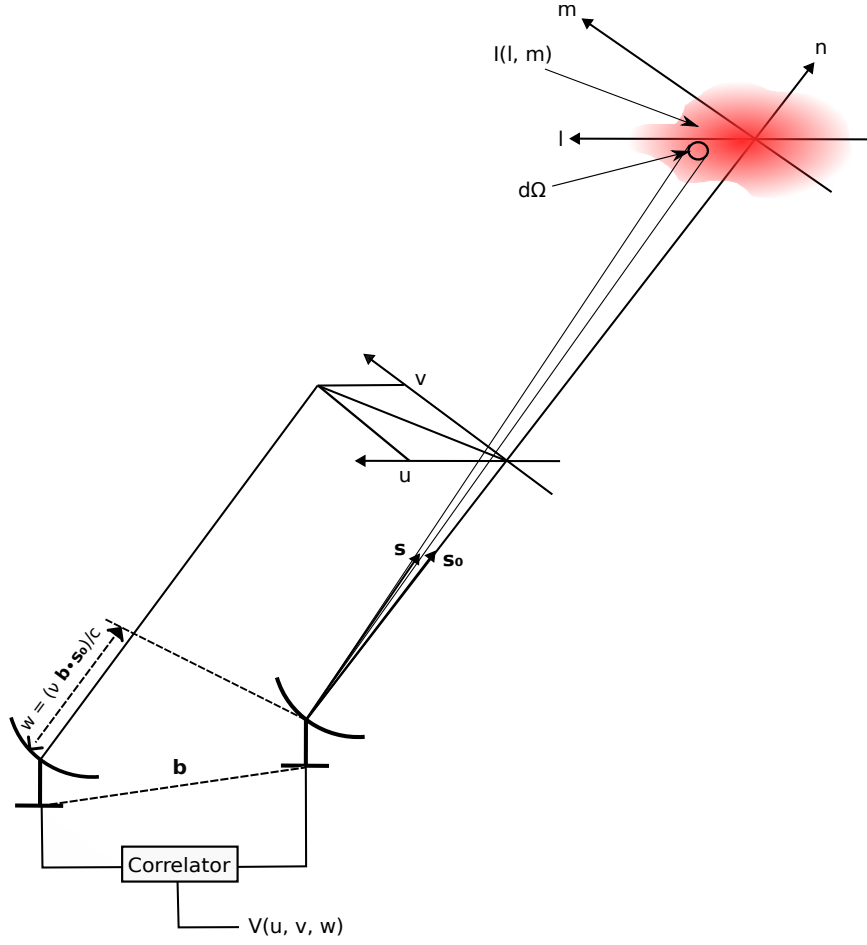


Figure 1.9: The (u, v, w) and (l, m, n) co-ordinate systems used to express the baselines and brightness distribution respectively.

and towards the phase tracking center $\mathbf{s}_0$ respectively. u , v and w are measured in wavelengths. The brightness distribution is described using the vectors (l, m, n) which are known as direction cosines and are measured with respect to the u, v and w vectors, with $n = \sqrt{1 - l^2 - m^2}$, as can be seen in Figure 1.9. By using this co-ordinate system we get:

$$\frac{\nu \mathbf{b} \cdot \mathbf{s}}{c} = ul + vm + wn \quad (1.35)$$

$$\frac{\nu \mathbf{b} \cdot \mathbf{s}_0}{c} = w \quad (1.36)$$

$$d\Omega = \frac{dl \, dm}{n} = \frac{dl \, dm}{\sqrt{1 - l^2 - m^2}} \quad (1.37)$$

The complex visibility equation then becomes:

$$V(u, v, w) = \int \int \mathcal{A}(l, m) I(l, m) e^{-2\pi i [ul + vm + w(\sqrt{1-l^2-m^2}-1)]} \frac{dl \, dm}{\sqrt{1-l^2-m^2}} \quad (1.38)$$

However, in the case of small-field imaging, l and m are small, and so the equation can be simplified and w can be neglected as follows:

$$\begin{aligned} \sqrt{1-l^2-m^2} &\approx 1 & w(\sqrt{1-l^2-m^2}-1) &\approx 0 \\ V(u, v) &= \int \int \mathcal{A}(l, m) I(l, m) e^{-2\pi i (ul + vm)} \, dl \, dm \end{aligned} \quad (1.39)$$

The complex visibility equation is now a 2-D Fourier transform of the brightness distribution, meaning that the brightness distribution of the source can be recovered by performing an inverse Fourier transform:

$$\mathcal{A}(l, m) I(l, m) = \int \int V(u, v) e^{2\pi i (ul + vm)} \, du \, dv \quad (1.40)$$

1.4.4 Calibration

Due to various environmental and instrumental effects, the visibilities measured by the interferometer will be corrupted. The relationship between the real visibilities V_{ij} and the measured visibilities V'_{ij} can be represented by the Hamaker-Bregman-Sault equation (Hamaker et al. 1996):

$$V'_{ij} = J_{ij} V_{ij} \quad (1.41)$$

where J_{ij} represents the effects corrupting the baseline between antennas i and j . When factored into its individual effects, this equation becomes:

$$V'_{ij} = B_i(\nu, t) B_j^*(\nu, t) g_i(t) g_j(t) V_{ij} e^{i(\theta_i(t) - \theta_j(t))} \quad (1.42)$$

B_i is the bandpass response for each antenna i . The bandpass response describes how the antenna gain varies as a function of frequency due to various frequency dependent

effects. g_i and θ_i are the amplitude and phase components respectively of the complex gain. The complex gain can vary with time, partly due to changes in the instrument but are mainly due to environmental and atmospheric effects. As the amplitude and phase fluctuations are uncorrelated, they can be solved separately. Another effect that needs to be accounted for is the *delay*, which is the difference in propagation time between the signals from different antennas. If not corrected for, a large gradient in phase will appear across the bandwidth. In order to correct for all of these effects, the data needs to be *calibrated*. This is usually done by observing a calibrator source, typically a bright source with a well-known visibility model. The calibration solutions are derived from this source and then applied to the target.

In addition to calibrating for time-dependent and frequency-dependent effects, the flux density scale of the data needs to be set since the correlator of the interferometer only measures signals in terms of their relative signal strength and their phase. Therefore it is necessary to observe a source with a well-known flux density which acts as the flux calibrator.

1.4.5 Low-Frequency Calibration

At low frequencies, such as with the LOFAR telescope, due to a combination of ionospheric effects and wide field-of-view, calibration becomes more complicated. At low frequencies, the ionosphere results in delays between the antennas causing phase errors. These delays are proportional to the electron column density along the line of sight through the ionosphere. Since this will vary looking through different parts of the ionosphere, with a large field-of-view this results in phase errors which vary within the observation field. Another issue with LOFAR is that since the stations are fixed and instead pointed by adjusting the delays between the antennas of the station, there are complex time-varying beams for each station, which are not well understood.

In order to solve for these direction-dependent effects (DDEs), several techniques have been developed such as the Factor pipeline (Van Weeren et al. 2016) or the DDF-Pipeline (Shimwell et al. 2019). These involve dividing up the observation field into

‘facets’, and then solving for the DDEs in each of these facets and applying the solutions to the data.

1.4.6 Imaging

Since the brightness distribution is given by the inverse Fourier transform of the visibilities, the visibility needs to be sampled at various points in the $u - v$ plane to get the brightness distribution. An interferometer array consisting of N antennas will contain $N(N - 1)/2$ pair combinations of antennas, each of which will sample a different part of the $u - v$ plane. As the earth rotates, the projected baseline of each antenna pair will also change and will trace out elliptical loci in the $u - v$ plane so different points in the $u - v$ plane will be sampled, this is known as *Earth-rotation aperture synthesis*. In addition, the $u - v$ plane can also be sampled by measuring visibilities over a range of frequencies. Since u and v are measured in units of wavelength, different values of u and v can be sampled without changing the baseline between the antennas. This technique is known as *multi-frequency synthesis*.

However as only a finite number of samples of the visibility function can be taken, $I(l, m)$ itself cannot be recovered directly. The set of points where the visibility is measured can be described by a sampling function $S(u, v)$, which is non-zero at points where the visibility is measured. The sampled visibility function is therefore given by $S(u, v)V(u, v)$. When taking the inverse Fourier transform of this, instead of obtaining the real image, one obtains what is known as the *dirty image*:

$$I^D(l, m) = \int \int S(u, v)V(u, v)e^{2\pi i(ul+vm)}dudv \quad (1.43)$$

Using the convolution theorem, it can be shown that the dirty image I^D is the real image I convolved with what is known as the *dirty beam* or *synthesized beam* (Taylor

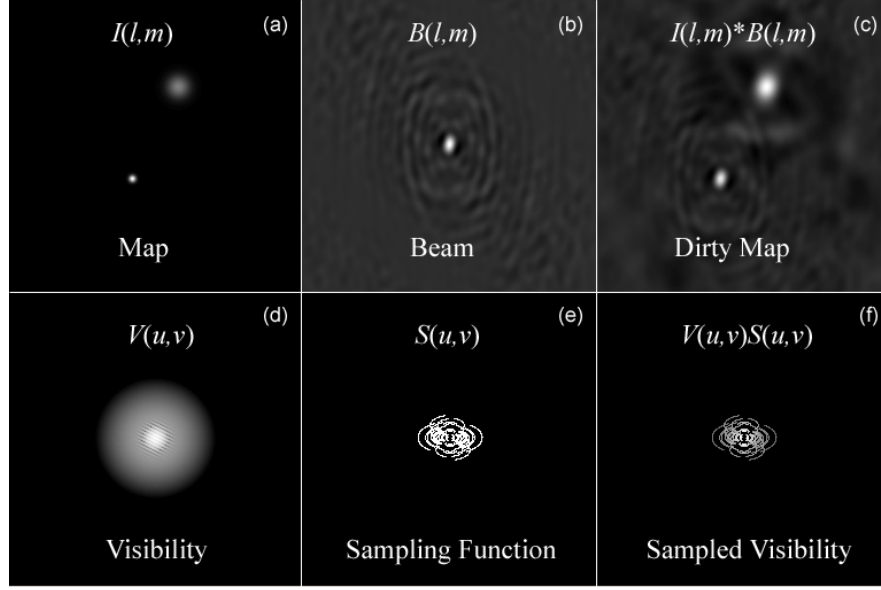


Figure 1.10: The Fourier transform relations for an example image (map). The upper row shows the: (a) sky map (b) dirty beam and (c) the resulting dirty map. The bottom row shows the corresponding: (d) visibilities (e) sampling function and (f) sampled visibilities. Credit: Dale Gary. Reproduced with permission of author.

et al. 1999):

$$I^D(l, m) = B^D(l, m) * I(l, m) \quad (1.44)$$

$$B^D(l, m) = \int \int S(u, v) e^{2\pi i(ul+vm)} du dv \quad (1.45)$$

The dirty beam $B^D(l, m)$ is simply the inverse Fourier transform of the sampling function. These relations are shown for an example brightness distribution in Figure 1.10.

To obtain the real image, I needs to be found by “deconvolving” the image. However, there is no unique solution for I as the points in the (u, v) plane where the visibility is not sampled can take any value and the convolution equation will still be fulfilled. The solution which has zero amplitude for all of the unsampled points is known as the “principal” solution. This is the solution which gives the dirty image. The principal solution is generally not a good solution for the image as it contains distortions known as sidelobes. In order to find the real image, reasonable values for the unsampled points in the (u, v) plane need to be found and combined with the principal solution.

1.4.7 CLEAN Algorithm

In order to deconvolve the image, the CLEAN algorithm (Högbom 1974) can be used. This is the most common technique for deconvolving radio images. It represents the radio image as series of point sources in an otherwise empty field and uses a simple iterative process to find the positions and strengths of these sources. The final image consists of these point sources convolved with a CLEAN beam, an elliptical Gaussian fitted to the central lobe of the dirty beam. The algorithm consists of the following steps:

1. The strength and position of the peak of the dirty image is found. If desired, this may be done only in certain areas of the image, known as CLEAN windows.
2. The dirty beam multiplied by the peak strength and a ‘gain’ factor γ (≤ 1), is subtracted from the dirty image at the position of the peak.
3. The position and strength of the peak is recorded and added to a model of the image.
4. The first 3 steps are repeated until any remaining peaks are below a user-specified level. The remainder of the image is now known as the *residuals*.
5. The final point source model is convolved with an idealised CLEAN beam (usually an elliptical Gaussian fitted to the central lobe of the dirty beam).
6. The residuals are added to the convolved point source model to create the final CLEAN image.

1.4.8 Weighting Functions

Before imaging is performed, the visibility data is usually weighted to control the shape of the synthesised beam Briggs (1995). Two commonly used weighting functions are *natural weighting* and *uniform weighting*.

Natural weighting weights all visibilities based simply on the inverse of the noise variance and so the weight of the i -th visibility is given by:

$$w_i = \omega_i = \frac{1}{\sigma_i^2} \quad (1.46)$$

where σ_i is the root-mean-square noise in the i -th visibility, and ω_i is the inverse variance in the i -th visibility. It tends to emphasise shorter baselines compared to uniform weighting, as there are more visibilities sampled at the centre of the (u, v) plane. It gives the best signal-to-noise ratio and so is good for detecting weak sources. However, it has worse angular resolution due to the emphasis on short baselines and also tends to have large sidelobe levels.

Uniform weighting weights visibilities inversely proportional to the density of sampled visibilities in the (u, v) plane, so it tends to emphasise longer baselines compared to natural weighting. First, the visibilities are gridded into cells in a regular, rectangular grid in the (u, v) plane. Then (u, v) the cells are re-weighted so that all of the cells have ‘uniform’ weights. The cell weights W_k are obtained by gridding the inverse variances ω_i on the (u, v) grid. The weight of the i -th visibility is then given by:

$$w_i = \frac{\omega_i}{W_k} \quad (1.47)$$

where W_k is the weight of the relevant cell k . This gives increased angular resolution and lower sidelobe levels but worse S/N ratio when compared to natural weighting.

There is also a hybrid form of uniform and natural weighting known as *robust weighting* or *Briggs weighting* Briggs (1995). This weighting scheme can provide a good trade-off between S/N ratio and sidelobe level. In this, the weights are given by:

$$w_i = \frac{\omega_i}{1 + W_k S^2} \quad (1.48)$$

where

$$S^2 = \frac{(5 \times 10^{-R})^2}{\frac{\sum_k W_k^2}{\sum_i w_i}} \quad (1.49)$$

The degree of trade-off is specified using the robustness parameter R which nominally varies between 2 and -2. For $R = 2$, S becomes very small and the weighting scheme approaches natural weighting. For $R = -2$, S becomes very large and the weighting scheme approaches uniform weighting.

1.4.9 Self-Calibration

After the calibration has been applied, there can still be residual phase and amplitude errors present in the data. If there is sufficient signal-to-noise ratio in the image, then these can be corrected using a technique known as self-calibration, where the image of the source is used as a model to calibrate the data, and thereby obtain a better image (Wilson et al. 2013). This can be done as the number of baselines is much larger than the number of antennas. Since the calibration is antenna-based, this means that the solutions are over-constrained.

After creating an initial image of the source using the CLEAN algorithm, a complex gain calibration is performed using a model of the target obtained from CLEAN. An improved image is then created using CLEAN, in order to obtain a better model of the source. Using the improved model from this image, another calibration is performed. This process is iterated until the self-calibration solutions converge. This can be done either for the phase only, or if the signal-to-noise is sufficient, it can be done for the amplitude and phase.

1.4.10 Polarisation and Stokes Parameters

Polarisation measurement is an important aspect of radio astronomy. Many types of radio emission show polarisation. For example, synchrotron emission typically shows a small degree of linear polarisation while electron-cyclotron maser emission is typically highly circularly polarised. Therefore, polarisation can help us determine the emission mechanism and determine properties of the source. In order to measure polarisation,

interferometers separate the emission into either its orthogonal linear components:

$$E_x(t) = E_{0_x} e^{i(2\pi\nu t + \delta_x)} \quad E_y(t) = E_{0_y} e^{i(2\pi\nu t + \delta_y)} \quad (1.50)$$

or circular components:

$$E_R(t) = E_{0_R} e^{i(2\pi\nu t + \delta_R)} \quad E_L(t) = E_{0_L} e^{i(2\pi\nu t + \delta_L)} \quad (1.51)$$

In radio astronomy, polarisation is usually measured in terms of the Stokes parameters (Wilson et al. 2013), which are defined in terms of either the linearly or circularly polarised components as:

$$I = \langle E_{0_x}^2 \rangle + \langle E_{0_y}^2 \rangle = \langle E_{0_R}^2 \rangle + \langle E_{0_L}^2 \rangle \quad (1.52)$$

$$Q = \langle E_{0_x}^2 \rangle - \langle E_{0_y}^2 \rangle = 2\langle E_{0_R} E_{0_L} \rangle \cos(\delta_R - \delta_L) \quad (1.53)$$

$$U = 2\langle E_{0_x} E_{0_y} \rangle \cos(\delta_y - \delta_x) = 2\langle E_{0_R} E_{0_L} \rangle \sin(\delta_R - \delta_L) \quad (1.54)$$

$$V = 2\langle E_{0_x} E_{0_y} \rangle \sin(\delta_y - \delta_x) = \langle E_{0_R}^2 \rangle - \langle E_{0_L}^2 \rangle \quad (1.55)$$

where the angular brackets represent the time averaged value. From these definitions, it can be seen that I is a measure of the total intensity of the emission, Q and U measure the linear polarisation, and V measures the circular polarisation. The linear and circular polarization fractions, m_l and m_c respectively, can be obtained from the Stokes parameters as follows:

$$m_l = \frac{\sqrt{Q^2 + U^2}}{I} \quad (1.56)$$

$$m_c = \frac{V}{I} \quad (1.57)$$

An important caveat to be aware of is that there are multiple conventions for defining the Stokes parameters within radio astronomy (Robishaw & Heiles 2021). For example, the radio astronomy software packages CASA and WSClean define Stokes I as the average of the two orthogonal components rather than the sum and so there

is an additional factor of $\frac{1}{2}$ in all of the Stokes parameters. There are also conflicting definitions regarding the sign of the Stokes V parameter. While the IAU defines Stokes V as being the right-hand circularly polarised component minus the left-hand circularly polarised component, many radio astronomy texts in the past have used the opposite definition i.e. the left-hand circularly polarised component minus the right-hand circularly polarised component. This is the definition typically used within pulsar astronomy for example. For each instrument, the sign of Stokes V is also affected by the location of the feed in the antenna (primary, secondary, tertiary focus). Therefore, it is necessary to check by comparing to sources with known circular polarisations (e.g. pulsars). This was done by Callingham et al. (2021b) for data from the LOFAR Two-Metre Sky Survey (LoTSS), who found that the data followed the pulsar convention. Since all of the LOFAR data in this survey is either from LoTSS or reduced and imaged using the same software as LoTSS, this is the convention used in this work for LOFAR data. Note that the convention used does not affect the polarisation fractions derived.

1.5 Instruments

1.5.1 LOFAR

The Low-Frequency Array (LOFAR) (Van Haarlem et al. 2013) is a low-frequency radio telescope array made up of 51 individual stations located mainly in the Netherlands and run by ASTRON, the Netherlands Institute for Radio Astronomy. There are 24 core stations spread out over a 3 km area located



near Exloo, in the northeastern Netherlands. Another 14 remote stations are spread out across the Netherlands. The core and re-

Figure 1.11: The LOFAR core in the Netherlands. Credit: LOFAR / ASTRON. Licensed under CC BY 3.0 (<https://creativecommons.org/licenses/by/3.0/>).

remote stations combine to give a maximum baseline of ~ 100 km. In addition to this, there are 13 international stations located across all of Europe including Germany, France, the United Kingdom, Sweden, Poland and Ireland. These give a maximum baseline of ~ 2000 km. Two other international stations are also currently planned for Latvia and Italy.

Each station consists of two types of antennas: the High-Band Antennas (HBAs) which cover 110 – 240 MHz and the Low-Band Antennas (LBAs) which cover 10 – 90 MHz. Unlike a traditional radio telescope, which consists of steerable antenna dishes, LOFAR contains no moving parts. Each station is instead composed of dozens of dipole antennas. The signals from these are combined digitally to form a phased array. Electronic beam-forming techniques are then used to “point” the telescope.

1.5.2 Giant Metrewave Radio Telescope (GMRT)



Figure 1.12: One of the GMRT antennas. Credit: Megan Argo. Licensed under CC BY 2.0 (<https://creativecommons.org/licenses/by/2.0/>)

The Giant Metrewave Radio Telescope (GMRT) (Swarup 1991) is a radio telescope array located 80 km north of Pune, India run by the Tata Institute for Fundamental Research (TIFR). It consists of 30 steerable antenna dishes, each of 45 m diameter, with a maximum baseline of 25 km, and observes at between 120 – 1460 MHz.. Recently, the GMRT underwent a major upgrade to install new, wideband receivers which provide much wider bandwidth and improved sensitivity.

1.5.3 Karl G. Jansky Very Large Array (VLA)

The Karl G. Jansky Very Large Array (VLA) is a radio astronomy observatory located fifty miles west of Socorro, New Mexico and operated by the National Radio Astronomy Observatory (NRAO). It is a radio interferometer consisting of 27 antennas in a Y-shape configuration, each with a diameter of 25 m. The VLA has 4 possible configurations: A, B, C and D with A being the largest configuration, with a maximum baseline of 36 km, and therefore



Figure 1.13: The VLA in its closest configuration, the D configuration. Credit: NRAO/AUI/NSF. Licensed under CC BY 3.0 (<https://creativecommons.org/licenses/by/3.0/>).

providing the highest resolution and D being the smallest configuration, and therefore

providing the lowest resolution. The frequency range is 54 MHz to 50 GHz. The array was first completed in 1980 and received a major upgrade in 2011 to provide greater frequency coverage and therefore much higher sensitivity.

1.5.4 e-MERLIN

The enhanced Multi Element Remotely Linked Interferometer Network (e-MERLIN) is a radio telescope array consisting of seven radio telescopes spread out across the UK. This includes the 76 m Lovell telescope, the 32 m Cambridge dish and the 25 m Mark II, Pickmere, Darnhall, Knockin, and Defford dishes. The maximum baseline of the array is 217 km, giving it a very high angular resolution, comparable to that of the Hubble Space Telescope. It operates at 3 radio bands: L band (1.23 - 1.74 GHz), C band (4.3 - 7.5 GHz), and K band (19 - 25 GHz).

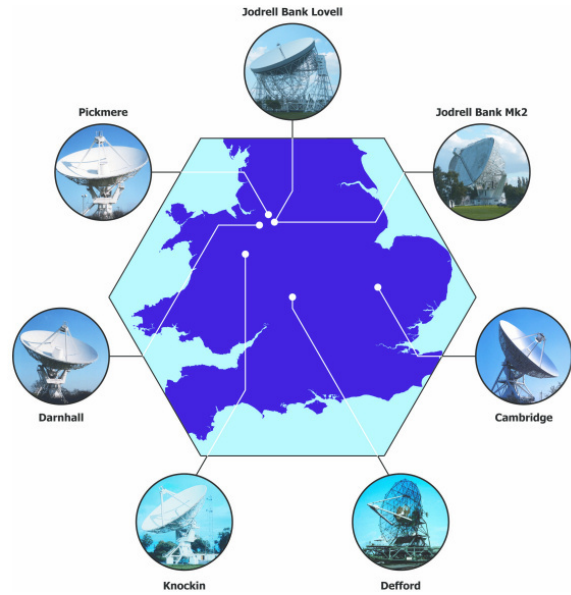


Figure 1.14: A map showing the locations of the e-MERLIN dishes across the UK. Credit: Alastair Gunn. Licensed under CC BY 3.0 (<https://creativecommons.org/licenses/by/3.0/>).

1.5.5 Atacama Large Millimeter/sub-millimeter Array (ALMA)



Figure 1.15: ALMA under the Milky Way. Credit: Sergio Otarola - ALMA (ESO/NAOJ/NRAO). Licensed under CC BY 4.0 (<https://creativecommons.org/licenses/by/4.0/>).

The Atacama Large Millimeter/sub-millimeter Array (ALMA) is a radio telescope located on the Chajnantor Plateau in the Atacama desert in Chile observing at wavelengths of 0.32 - 3.6 mm (80 - 940 GHz). It is run by an international collaboration consisting of the European Southern Observatory (ESO), the NRAO, Canada, Japan, South Korea, Taiwan, and Chile.

It consists of 66 antennas, including the main array of 50 12-m antennas and a

compact array of 4 12-m antennas and 12 7-m antennas. These antennas can be arranged in different configurations, giving baselines varying from 150 m to 16 km, resulting in very high resolution at these wavelengths.

1.6 Thesis Outline

As shown in this chapter, radio astronomy offers many opportunities to learn more about the process of star formation. In this thesis, I will show exciting new ways of using radio astronomy to further our understanding of YSOs which I have applied in my work. In the next two chapters in particular, I will discuss how low-frequency observations of YSOs using instruments such as LOFAR, an area which has not been explored much previously, can help us.

In chapter 2, I report on the detection of a low-frequency spectrum turnover in an emission knot within the jet of the low-mass YSO DG Tau A using LOFAR, the first time such a turnover has been detected in nonthermal emission in a YSO jet. I describe the possible mechanisms that could explain this turnover, one of which, the Razin effect, could provide an exciting new method of studying the magnetic field strength within YSO jets.

In chapter 3, I report on the detection of coherent radio bursts from two YSOs. I discuss the two possible mechanisms to generate the emission and why it is most likely to be ECM emission. From this, I estimate the magnetic field strength in the source region and describe how the ECM emission may be generated. One explanation, a star-planet interaction, is particularly interesting as it may provide a method to infer the presence of exoplanets around YSOs.

Then in chapter 4, I use high-resolution radio observations obtained using the VLA, e-MERLIN, and ALMA to study the jets from the binary YSOs L1551 IRS 5 and L1551 NE. By comparing these observations with previous ones, it is shown that there is significant variability in the flux density of the free-free emission of the jets of these objects. From this, it is also inferred that there is variability in the ionised mass loss rate of the jets. As well as this, radiative transfer modelling is used to obtain a model image of one of the jets. By comparing this with the observed image, an estimate for the collimation distance is obtained.

Finally, in chapter 5, the overall conclusions of this thesis are presented and suggestions for future work are discussed.

Chapter 2

The first detection of a low-frequency turnover in nonthermal emission from the jet of a young star

Young stars are associated with powerful jets which carry away mass and angular momentum from the system (Frank et al. 2014). Though the exact collimation and launching mechanism for the jets is uncertain, it is generally agreed that it involves a magnetic field originating in the star or the accretion disk (Pudritz & Norman 1983b; Shu et al. 1994). Material from the accretion disk is lifted and centrifugally accelerated along the magnetic field lines. The jets are then also collimated by the magnetic field due to “hoop stresses” from the toroidal component of the field. However, despite their importance to the origin of the jet, so far magnetic fields in jets are poorly understood (Ray 2009) and measurements of the field strength are scarce.

Radio emission from YSOs has generally been observed to have spectral indices in the range $-0.1 < \alpha < 1$ (Anglada 1996), where the flux density S_ν at frequency ν , $S_\nu \propto \nu^\alpha$. This is characteristic of thermal bremsstrahlung emission and indicative of partially ionized outflows (Reynolds 1986), where the bulk of the emission comes from

the base of the ionized jet.

In contrast, nonthermal emission from YSO jets, characterised by negative spectral indices, is not very well studied even though it could provide valuable information about the poorly studied magnetic field (Ray 2009). This is mainly because of the high sensitivities required to observe the normally weak nonthermal emission. Recently however, nonthermal emission has been observed towards several YSOs, generally in radio knots in the jet, which show negative spectral indices (Anglada et al. 2018). Its origin is thought to be synchrotron emission from relativistic electrons, accelerated either in shocks caused by the collision of the material within the jet with the dense gas in the surrounding molecular cloud, or in internal shocks within the jet due to variable ejection rates (Padovani et al. 2015, 2016). Particle acceleration is achieved through the process of diffusive shock acceleration (DSA), a first-order Fermi mechanism (Drury 1983). Recently, Purser et al. (2016) found that out of 28 high-mass ($M > 8 M_{\odot}$) YSOs associated with jets, 14 of them were also associated with nonthermal radio emission, suggesting that this could be a common feature of high-mass YSO jets. In the high-mass YSO jet associated with the Herbig Haro objects HH80 and HH81, Carrasco-González et al. (2010) detected linear polarisation in the jet, showing the presence of synchrotron emission and the orientation of the magnetic field.

DG Tau A is a classical T-Tauri star located in the Taurus Molecular Cloud at a distance of 120.8 ± 2.2 pc, based on GAIA data (Gaia Collaboration et. al 2016; Bailer-Jones et al. 2018). It is associated with a bipolar outflow and was one of the first YSOs to be associated with an optical jet (Mundt & Fried 1983), displaying several shocks and knots along its axis. The optical jet has a position angle of 223° (Lavalley et al. 1997) and an inclination to the line of sight of $37.7^{\circ} \pm 2.2^{\circ}$ (Eisloffel & Mundt 1998). In the radio, close to the source, the emission is elongated in the direction of the outflow and has a spectral index typical of thermal free-free radiation (Lynch et al. 2013).

Further from the source, the outflow exhibits synchrotron emission. In 2012, it was observed using the Giant Metrewave Radio Telescope (GMRT) at 323 MHz and 608 MHz (Ainsworth et al. 2014, 2016). At both frequencies, an extended region of

nonthermal emission, hereafter referred to as knot C, was detected $\sim 14''$ southwest of DG Tau A or ~ 3000 au, slightly offset from the axis of the outflow. This was hypothesised to be associated with the bow shock of the jet. Combined with previous observations from the Karl G. Jansky Very Large Array (VLA) at 5.5 GHz and 8 GHz (Lynch et al. 2013), a spectral index of -0.89 was measured. Assuming equipartition between the energy of the relativistic electrons and protons in the source and the magnetic field energy (Beck & Krause 2005b), the magnetic field strength was estimated to be 110 μ G. Another region of nonthermal emission, hereafter referred to as knot D, was also detected to the northeast of DG Tau A at 323 MHz (Ainsworth et al. 2016), which was assumed to be associated with the counter-jet.

These two emission knots were later detected again at 6 GHz and 10 GHz using the VLA by Purser et al. (2018). No proper motion was detected in knot C over 4 years suggesting it may not be associated with a bow shock but instead with a quasi-stationary shock due to dense cloud material drifting into the jet. This could also be the case for knot D.

In this chapter, I report on the observation of DG Tau A at 152 MHz using the Low-Frequency Array (LOFAR) (Van Haarlem et al. 2013). This is the second time that a YSO has been observed using LOFAR. The low-mass YSO T Tau was observed in the same observing campaign and reported on in Coughlan et al. (2017). In section 2.1, I discuss how the observations were carried out and the data reduction process. In section 2.2, I describe the resulting image obtained. In section 2.3, I then describe the modelling of the low-frequency turnover observed in the spectrum of knot C in the jet of DG Tau A. Finally, I present the concluding remarks of this chapter in section 2.4.

2.1 Observations and Data Reduction

DG Tau A was observed on 2013 November 30 - December 1 (epoch 2013.91), using LOFAR (Project code: LC1.001). Total on source time was 8 hours with the High Band Array (HBA) using 74 MHz of bandwidth, and a central frequency of 152 MHz.

Before I received the data, the data were run through the pre-processing pipeline by

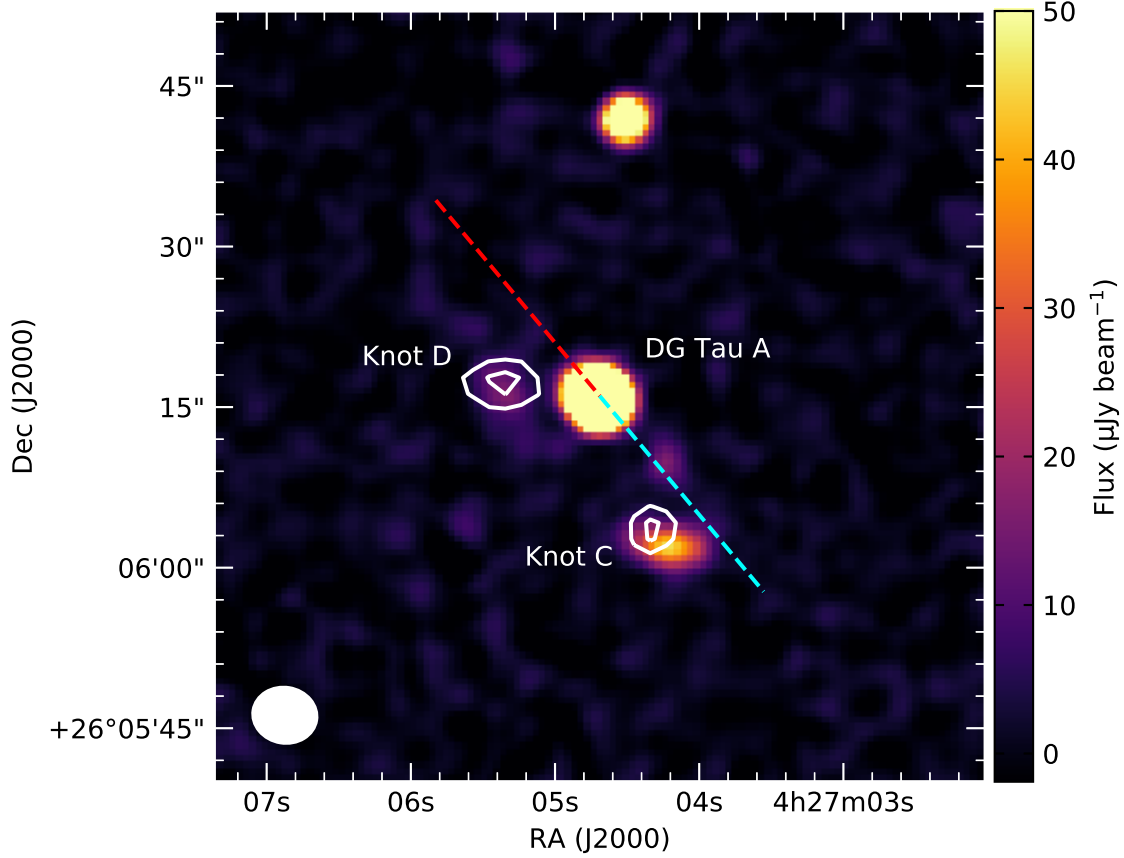


Figure 2.1: Contour plot of the DG Tau A observation at 152 MHz overlaid on a colour map of a 6 GHz image of DG Tau A from the VLA (Purser et al. 2018). The restoring beam for the LOFAR observation was $6.04'' \times 5.25''$ with a position angle of 80.7° and is shown in the bottom left corner. The root-mean-square noise level in the LOFAR observation was $\sigma_{rms} = 90 \mu\text{Jy beam}^{-1}$ and the contour levels are $-3, 3$ and $4 \times \sigma_{rms}$. The noise level in the VLA observation was $\sigma_{rms} = 1.9 \mu\text{Jy beam}^{-1}$. The receding and approaching lobes of the jet are indicated by the red and blue dashed lines respectively.

Epoch	Instrument	ν (MHz)	σ_{rms} ($\mu\text{Jy beam}^{-1}$)	Knot C (μJy)	Knot D (μJy)
2013.91	LOFAR	152	90	490 ± 145	860 ± 205
2012.95	GMRT	323	127	1350 ± 207	936 ± 196
2012.95	GMRT	608	80	1087 ± 175	–
2015.10	VLA	2500	8	219 ± 29	122 ± 34
2015.10	VLA	3500	5	159 ± 23	–
2012.22	VLA	5400	6.2	73 ± 8	–
2016.15	VLA	6000	1.9	95 ± 6	42 ± 4
2012.29	VLA	8500	8	39 ± 9	–
2016.10	VLA	10000	2.7	72 ± 15	–

Table 2.1: Knot C and knot D flux density measurements

the ASTRON Radio Observatory. This automatically flagged the data for RFI using AOFLAGGER (Offringa et al. 2012) and then averaged in time to 4 s and in frequency to 50 kHz per channel. For LOFAR, there are important direction-dependent effects (DDEs) present due to the ionosphere and imperfect knowledge of the beam shapes, therefore a technique known as facet calibration (Van Weeren et al. 2016) is required to accurately calibrate the data.

Initially, direction-independent calibration was performed using the LOFAR pre-facet calibration pipeline, known as PREFACTOR. Amplitude and phase gain calibration was performed using the calibrator source 3C147 using the flux density scale from Scaife & Heald (2012). This source was observed in two 10 minute runs, bookending the observations of the target field. The phase solutions were then separated into the contributions due to clock offsets of the individual stations and due to the total electron content (TEC) of the atmosphere. Both the amplitude and clock solutions were transferred from the calibrator to the target field. However, the TEC solutions were not transferred as these are direction-dependent and so they are not applicable to the target field. PREFACTOR then performed an initial direction-independent phase calibration using a model of the target field obtained from the TIFR GMRT Sky Survey (TGSS) (Intema et al. 2017). Using WSCLEAN (Offringa et al. 2014), each 2 MHz band was then imaged and the clean models produced were then subtracted from the visibilities and converted into sky models. This results in a series of 2 MHz bands with all the sources subtracted and a corresponding sky model for each band.

To account for the DDEs, the LOFAR facet calibration pipeline, known as FACTOR, was used. This divides the field up into ‘facets’ each of which is centred on a bright, compact source or group of sources which acts as a calibrator. The calibrators are selected based on their brightness and angular size. A brightness threshold of 0.1 Jy and a maximum size of $2'$ was used to select the individual sources. Sources which had a separation of less than $9'$ were grouped together as a calibrator group. Sources, or groups of sources, whose total combined flux density was brighter than 0.2 Jy were then used as the calibrators. For each facet, the calibrator sources were added back to the data and the data were phase shifted to the position of the calibrator field. The calibrator field was then imaged and several rounds of self-calibration were performed, until the noise was no longer significantly improving. The other sources from the sky model in the entire facet were then added back to the data and the facet was imaged with the solutions for the calibrator field applied to the whole facet. After imaging the facet, an improved sky model was extracted from the image and subtracted from the data with the improved DDE solutions applied. This process was repeated for all of the facets in order of decreasing brightness of the calibrator, with the exception of the facet containing the target which was processed last to ensure that all the artifacts from other facets were removed before the target facet was calibrated.

The bright outlier source 3C123 was also subtracted from the data using the *outlierpeel* task in FACTOR. A model of 3C123 was obtained using an observation from the LOFAR Long-Term Archive (LTA). The data from this observation were calibrated using PREFACTOR and then imaged using WSCLEAN. A model of 3C123 was then extracted from this image using the PYBDSF (Mohan & Rafferty 2015) source extraction software and used to calibrate and subtract the source from the DG Tau A observation.

Parts of the frequency spectrum were flagged due to poor amplitude solutions found by PREFACTOR, likely due to RFI. The final half hour of the observation on all stations, and the entire last hour on station RS407HBA, was also flagged due to poor calibration solutions found by FACTOR.

After the calibration was complete, all of the sources for each facet were added back to the data and each facet was re-imaged. Final imaging was done using the Briggs weighting scheme (Briggs 1995) with a robust parameter of -0.5.

Due to inaccuracies in the LOFAR HBA beam model (Van Weeren et al. 2016; Shimwell et al. 2016), there are large uncertainties in the flux scale. To account for this, the program TOPCAT (Taylor 2005) was used to compare the integrated flux densities of several compact bright sources in the field of the LOFAR image with their flux densities in the TGSS survey. The average ratio of TGSS flux densities to LOFAR flux densities was found to be 0.83 ± 0.12 . Therefore, the flux densities measured in the LOFAR image were corrected by this factor to account for this. An absolute flux density calibration error for LOFAR of 15% was then assumed for all the measured flux densities, as used in other LOFAR HBA observations (Shimwell et al. 2016; Savini et al. 2018).

Due to the ionosphere there can be large astrometric errors present in LOFAR images. Therefore, the positions of the bright compact sources in the field were compared with their positions in the TGSS survey. A relatively large systematic offset was found of $1.88'' \pm 1.50''$ with a position angle for the offset of $11^\circ \pm 54^\circ$.

2.2 Results

After calibrating the visibilities and reducing the data, a radio image of DG Tau A was produced (see Figure 2.1).

To measure the flux densities in the image, the task *imfit* in the Common Astronomy Software Application (CASA) (McMullin et al. 2007) was used. This measures the integrated flux density by fitting a Gaussian to the emission. The error in the flux density measurements was taken to be a combination of the root-mean-square noise σ_{rms} of the image around DG Tau A, the fitting error σ_{fit} from the Gaussian fit to the emission and the absolute flux calibration error, which was 15% for LOFAR:

$$\sigma_{S_\nu} = \sqrt{\sigma_{rms}^2 + \sigma_{fit}^2 + (0.15 \times S_\nu)^2}.$$

Both knot C and knot D were detected with flux densities of $490 \pm 145 \mu\text{Jy}$ and

$860 \pm 205 \mu\text{Jy}$ respectively. These measurements were combined with previous GMRT and VLA measurements (Ainsworth et al. 2016; Lynch et al. 2013; Purser et al. 2018) and VLA archival data to give the spectra in Figure 2.2 and are also listed in Table 2.1. Unfortunately, the thermal jet was not detected as it is too weak at these frequencies, although a 3σ upper limit of $\lesssim 270 \mu\text{Jy}$ can be placed on its peak flux density.

The image in Figure 2.1 appears to show a difference in the position of the peak of the emission for knot C between the LOFAR image and the VLA image at 6 GHz. The peak position in the LOFAR image was offset compared to the VLA image by $2.39''$ with a position angle of 31° . This is within the limits of the systematic offset between the LOFAR image and the TGSS survey suggesting that the difference in position is most likely an astrometric error due to the systematic offset in the LOFAR image.

The integrated flux densities for the GMRT observations (Ainsworth et al. 2014, 2016) at 323 MHz and 608 MHz were remeasured using *imfit* so that a consistent method to measure the flux densities was used for comparison purposes. At 608 MHz, due to the curved shape of the emission at this frequency, it was necessary to use two Gaussians to accurately model the flux density. To ensure that this approach was accurate, the model was subtracted from the image. The residual image was then inspected to check that essentially all the emission from the source had been removed. The errors were calculated with the same procedure as for the LOFAR measurements, except with a flux density calibration error of 5%.

In the case of knot C, the emission is clearly much lower than predicted for non-thermal emission (see Figure 2.2). Only limited flux density variability was found at higher frequencies, where there was a $\sim 2\sigma$ increase in flux density at 5.5 GHz and 8.5 GHz between 2012 and 2016 (Purser et al. 2018), making it unlikely that there was a significant decrease in flux density between observations. This suggests that a spectral low-frequency turnover has been observed.

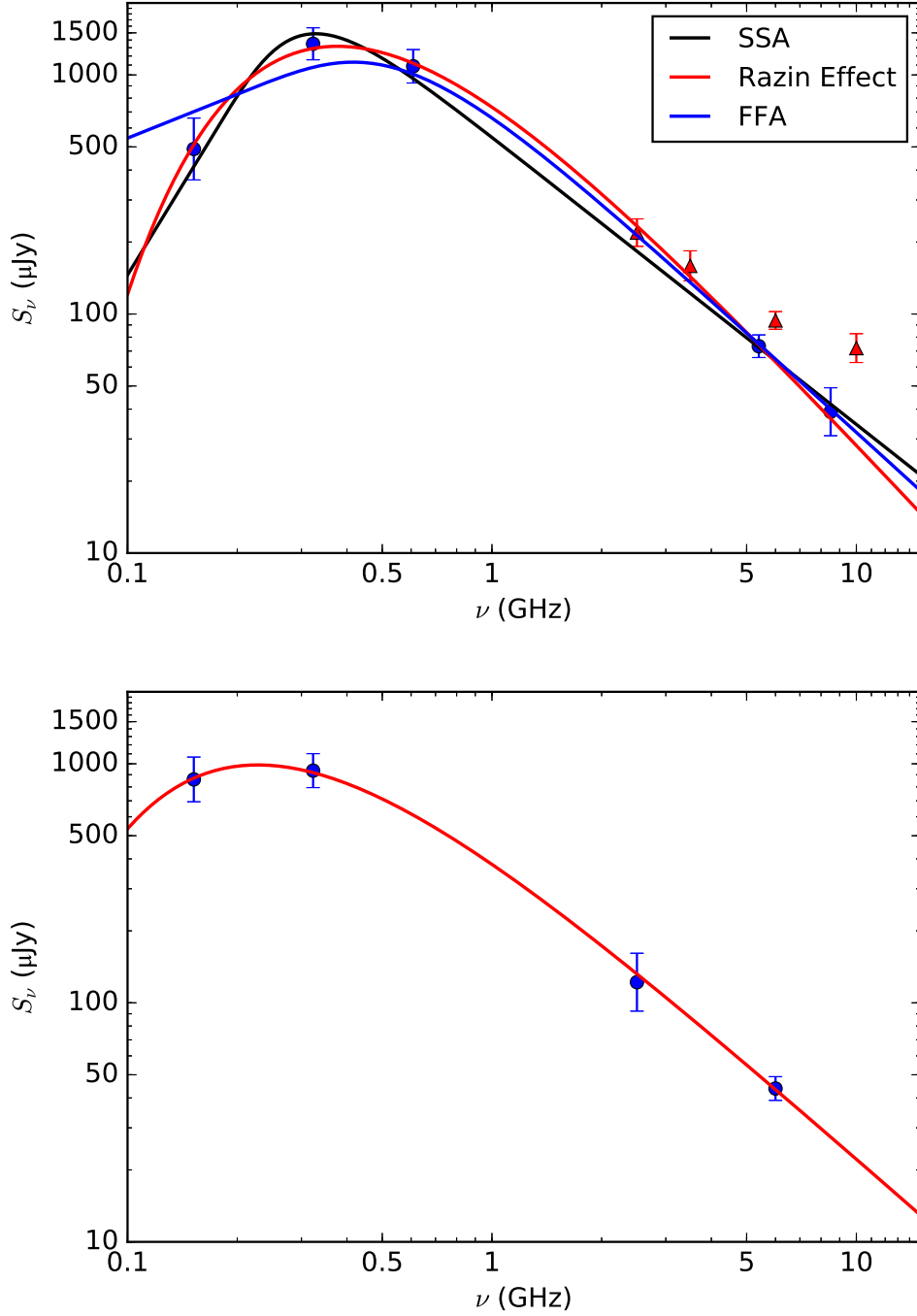


Figure 2.2: **Top:** Flux density values of knot C plotted against frequency. Different turnover mechanisms were fitted to the data from LOFAR at 152 MHz, GMRT at 323 MHz and 608 MHz (Ainsworth et al. 2014) and the 2012 VLA observations at 5.5 GHz and 8.5 GHz (Lynch et al. 2013), which are indicated by the blue circles. Flux densities from 2015 VLA archival data at 2.5 GHz and 3.5 GHz and from previously published 2016 VLA data at 6 GHz and 10 GHz (Purser et al. 2018) are also plotted, indicated by the red triangles, However, these were not fitted as they are from 2015 and 2016 and too different in time to be comparable. **Bottom:** Flux density values of knot D plotted against frequency. Data were from LOFAR at 152 MHz, GMRT at 323 MHz (Ainsworth et al. 2014), 2015 VLA archival data at 2.5 GHz and previously published 2016 VLA data at 6 GHz (Purser et al. 2018). A Razin effect turnover has been fitted to the data.

2.3 Discussion

2.3.1 Finding the Low-Frequency Turnover Mechanism of Knot C

To find the low-frequency turnover mechanism responsible, models for each potential mechanism were tested and fitted to the spectrum. Only the measurements from 2012 and 2013 were considered in order to avoid any issues with flux variability with time. An isotropic, homogenous, spherically symmetric, emitting region was assumed for each of the models with a uniform, randomly oriented magnetic field.

The flux density produced can be obtained from the equation of radiative transfer:

$$S(\nu) = \Omega \frac{j_\nu}{\alpha_\nu} (1 - e^{-\alpha_\nu l}) \quad (2.1)$$

where Ω is the solid angle of the emitting region, j_ν is the emissivity coefficient, α_ν is the absorption coefficient and l is the size of the emitting region.

At first, the most obvious explanation would appear to be synchrotron self-absorption (SSA) (Rybicki & Lightman 1979). Since the physical parameters in the model are coupled, combined parameters were fitted to the spectrum, based on Equation 1.17 and Equation 1.18. The parameters fitted were p , P_1 and P_2 , where $P_1 \equiv \Omega B^{-1/2}$ and $P_2 \equiv \kappa B^{(p+2)/2} l$. From the value for the combined parameter P_1 , calculated from the spectrum fit, the magnetic field strength required in the source for SSA to occur can be calculated if the angular size of the source is known:

$$B = \left(\frac{\Omega}{P_1} \right)^2 \quad (2.2)$$

While knot C is unresolved at 150 MHz, it was resolved by Purser et al. (2018) at 6 GHz and 10 GHz using the VLA, and found to have a deconvolved size of $5.3'' \times 1.3''$. Therefore I used this as an estimate for the size of the emitting region. SSA seems to provide a good fit to the data and gives a turnover frequency of 299 MHz. However, in order to fit the spectrum, it would require a magnetic field strength in the knot of

$B = 2.5 \times 10^{12}$ G, a highly unrealistic value and in complete disagreement with the equipartition value previously derived ($B = 110 \mu\text{G}$).

It should be noted that the derived magnetic field strength depends strongly on the angular size of the emission region ($B \propto \theta^4$). However, it would need to be many orders of magnitude smaller to give realistic values for B . To obtain the equipartition magnetic field strength for example, with the derived value of P_1 , the angular size would need to be $\sim 2 \times 10^{-4}$ arcsec. At the distance of DG Tau A, this corresponds to a size of ~ 0.025 au. This seems unlikely given that at higher frequencies, knot C has been resolved and found to have a deconvolved size of $5.3'' \times 1.3''$ (Purser et al. 2018). In addition, the emission is located at $14''$ or ~ 3000 au along the jet from the central star. Given the opening angle of the jet, which was estimated by (Maurri et al. 2014) to be $\sim 16 - 38^\circ$, this would mean that the jet is $\sim 800 - 2000$ au wide at the location of the emission region. This would require a very unusual geometry where the accelerating shock and resulting synchrotron emission is concentrated in a very small portion of the jet. Therefore, another explanation is clearly required.

Free-free absorption (FFA) (Rybicki & Lightman 1979) in the emission knot could also provide an explanation. Again, combined parameters were used when fitting the spectrum, based on Equation 1.17 and Equation 1.12. The parameters fitted were p , P_3 and P_4 , where $P_3 \equiv \kappa \Omega B^{(p+1)/2} T^{3/2} n_e^{-2}$ and $P_4 \equiv T^{-3/2} n_e^2 l$. From the combined parameter P_4 , the electron density in the source required for FFA to occur can be calculated if T and l are known:

$$n_e = \sqrt{\frac{P_4 T^{3/2}}{l}} \quad (2.3)$$

From optical emission line ratio observations (Oh et al. 2015), I assume $T_e \sim 5000$ K and the size of the emitting region was again taken to be the beam size as an upper limit ($\theta \approx 5''$). However, in order to fit the spectrum, an electron density of $n_e \approx 1.2 \times 10^4 \text{ cm}^{-3}$ would be required, much higher than the values seen from emission line ratios of $n_e \sim 500 \text{ cm}^{-3}$ (Oh et al. 2015). If the region is smaller than assumed, this would require an even higher electron density. Given this, FFA in the emission

knot seems unlikely. While FFA along the line of sight is also possible, this also seems improbable given that knot D exhibits a significantly shallower low-frequency turnover, and so this would require the column density along the line of sight to vary between these two regions.

Another possible explanation is a low-energy cut-off in the energy distribution of the electrons resulting in a low-frequency cut-off in the radiation spectrum. Below the frequency cut-off, the radiation will have a spectral index of $+1/3$ or less, since this is the low-frequency spectral index of radiation from a single electron (Rybicki & Lightman 1979) so the spectral index cannot be steeper than this. However, this would not provide a good fit to the data, since the spectral index below the low-frequency turnover is much steeper than $+1/3$.

Finally, the low-frequency turnover could be due to the Razin effect, also known as the Razin-Tsytoich effect (Tsytoich 1951; Razin 1960). For synchrotron radiation, the emitting source is a plasma, and so the refractive index, n_r , is less than 1. Therefore, the phase velocity of light in the plasma becomes c/n_r , with n_r decreasing towards low frequencies. This suppresses the beaming effect responsible for synchrotron radiation at low frequencies and therefore the synchrotron emission is reduced.

For the Razin effect, which is a reduction in the emission as opposed to an absorption effect like FFA or SSA, the low-frequency turnover has a form where below the turnover frequency, or Razin frequency, ν_R , the flux density decreases exponentially with frequency and it has a noticeable effect above ν_R , e.g. at $10\nu_R$ there is still a 10% reduction in flux (Hornby & Williams 1966). Therefore, the low-frequency turnover can be approximated by multiplying the optically thin spectrum by a factor of $e^{-\nu_R/\nu}$ (Dougherty et al. 2003). The parameters fitted are a constant of proportionality K , ν_R and p .

The turnover frequency is given by Ginzburg & Syrovatskii (1965):

$$\nu_R = 20 \frac{n_e}{B} \text{ MHz} \quad (2.4)$$

where n_e is the electron density in cm^{-3} and B is the magnetic field strength in μG . The

Razin effect was found to provide a good fit to the data and gives a turnover frequency of $\nu_R = 629 \pm 30$ MHz. From the S II emission line ratio (Oh et al. 2015), the electron density in the jet at $14''$ from the source is known to be $n_e \sim 500 \text{ cm}^{-3}$. Based on this, and the turnover frequency derived, a magnetic field strength of $B \sim 20 \text{ } \mu\text{G}$ was calculated.

The constant of proportionality K is related to the constant κ in the electron energy distribution in Equation 1.16 by the equation:

$$K = \frac{2.344 \times 10^{-25}}{4\pi} (1.253 \times 10^{37})^{(p-1)/2} a(p) \Omega l B^{(p+1)/2} \kappa \quad (2.5)$$

where $a(p)$ is a constant dependent on p , obtained from Longair (2011). If κ is obtained from the value for K for the fit to the spectrum, an estimate for the density of relativistic electrons can then be obtained by integrating over the range of relevant electron energies. The energy E of an electron whose peak emission intensity is at frequency ν is given by the equation:

$$E = \left(\frac{\nu}{CB} \right)^{1/2} \quad (2.6)$$

where $C = 1.22 \times 10^{10} / (m_e c^2)^2$. Therefore, the minimum and maximum electron energies can be taken to be the energies corresponding to the minimum and maximum frequencies observed. For the Razin effect fit, for an angular size of $5''$ and using the values of B and the electron power law index p derived from the fit, it was found to be $n_e^{\text{rel}} \sim 5 \times 10^{-6} \text{ cm}^{-3}$.

Based on all the possibilities considered, the Razin effect appears to be the most likely explanation for the low-frequency turnover, as it is the only explanation which provides a realistic spectral fit and plausible physical parameters. It should be noted that regardless of whether the Razin effect is responsible, this calculation still provides a lower limit on the magnetic field strength as if it were lower, the spectrum would already have turned over at a higher frequency due to the Razin effect.

The fitting parameters for all of the turnover mechanisms fitted are listed in Ta-

SSA:			
P_1 ($T^{-1/2}$)	P_2 ($J^{p-1} \text{ m}^{-2} T^{(p+2)/2}$)	p	α
$(8.06 \pm 2.01) \times 10^{-15}$	$(2.03 \pm 7.45) \times 10^{-23}$	3.41 ± 0.12	-1.20 ± 0.06
FFA:			
P_3 ($J^{p-1} \text{ m}^3 T^{(p+1)/2} K^{3/2}$)	P_4 ($K^{-3/2} \text{ m}^7$)	p	α
$(4.08 \pm 77.68) \times 10^{-71}$	$(3.63 \pm 3.84) \times 10^{-28}$	3.78 ± 0.57	-1.38 ± 0.29
Razin effect:			
K ($\text{W m}^{-2} \text{ Hz}^{(p-3)/2}$)	ν_R (MHz)	p	α
$(1.22 \pm 1.10) \times 10^{18}$	625 ± 32	4.32 ± 0.08	-1.66 ± 0.04

Table 2.2: Fitting parameters for each of the turnover mechanisms fitted to knot C of DG Tau A. Note that the units for some of the parameters depend on the electron power law index p . The optically thin spectral indices α corresponding to the values of p derived are also listed.

ble 2.2, as well as the corresponding values of the optically thin spectral index α , which is related to p by the relation $\alpha = -(p - 1)/2$. The values of α calculated for each of the fits were -1.20 ± 0.06 , -1.39 ± 0.29 and -1.66 ± 0.04 for SSA, FFA and the Razin effect respectively. These values for α are significantly steeper than the value measured by Ainsworth et al. (2014), which was -0.89 ± 0.07 , and also significantly steeper than the maximum value expected for a synchrotron source ($\alpha \approx -1$), unless spectral ageing is involved. However, it should be noted that there are only two points at higher frequencies (5.5 GHz and 8.5 GHz), which are quite close in frequency space, meaning that it is hard to be certain about the accuracy of the estimates for the spectral index of the optically thin region of the spectrum. The data points for the 2015 and 2016 VLA data seem to show a shallower spectral index, but as mentioned previously these were not included in the fit to avoid potential issues with variability.

2.3.2 Testing spectrum fits

Each of the fits for the low-frequency turnover mechanisms fitted to the spectrum of knot C was tested using the reduced chi-squared test to check the accuracy of the fit.

This is given by:

$$\chi^2_\nu = \frac{1}{k} \sum_{i=1}^N \frac{(y_i - f(x_i))^2}{\sigma_i^2} \quad (2.7)$$

where y_i are the measured data values, $f(x_i)$ are the theoretical data values based on the parameters of the fit, σ_i are the errors for the data values and k is the degrees of freedom which is given by $k = n - m$ where n is the number of data values and m is the number of fitted parameters. A value of $\chi^2_\nu \approx 1$ indicates a good fit while $\chi^2_\nu \gg 1$ indicates a poor fit. The values calculated for each fit are listed in Table 2.3.

Turnover Mechanism	χ^2_ν
SSA	0.65
FFA	2.05
Razin effect	0.11

Table 2.3: χ^2_ν values for each of the fitted low-frequency turnover mechanisms

2.3.3 Comparing with Previous B Estimates

Despite the importance of the magnetic field to the jet launching mechanism and collimation, gauging the strength of the magnetic field in YSO jets has proven very difficult. As a result, measurements of the magnetic field strength are scarce. Therefore, if the Razin effect is responsible for the turnover, this would provide a valuable measurement of the magnetic field strength along a YSO jet. In optical jets, the pre-shock magnetic field strength has previously been estimated through its effect on the post-shock compression (the ratio of the post-shock electron density to the pre-shock electron density). For the distant bow shocks in the jets of two less evolved YSOs (Morse et al. 1992, 1993), which are more embedded within their surrounding envelopes, B was estimated to be $\sim 20 - 30 \mu\text{G}$ for $n_e \sim 100 - 200 \text{ cm}^{-3}$ at distances of $\sim 5 - 6 \times 10^4 \text{ au}$ from the YSO. This compares well with the measurement for DG Tau A of $\sim 20 \mu\text{G}$ at $\sim 3 \times 10^3 \text{ au}$ from the source. The previously mentioned HH80-81 jet by comparison, had a larger field strength of $\sim 200 \mu\text{G}$, $\sim 1 \times 10^5 \text{ au}$ from the source, measured using the equipartition method, although this is for a massive YSO, so the conditions in the jet may be very different.

2.3.4 Equipartition Magnetic Field Strength

The equipartition magnetic field was also recalculated based on the spectrum observed in this work. The standard formula for the equipartition magnetic field strength is based on the magnetic field corresponding to the minimum total energy in the source required to generate the synchrotron radiation observed and is given by Longair (2011) as:

$$B_{eq} = \left[\frac{3\mu_0}{2} \frac{G(\alpha)(1+k)L_\nu}{Vf} \right]^{2/7} \quad (2.8)$$

where μ_0 is the vacuum permeability, $G(\alpha)$ is a constant dependent on α and the minimum and maximum observed frequencies of the radio spectrum, k is the ratio of the energy of relativistic protons in the source to that of relativistic electrons in the source (typically ≈ 40 for electrons undergoing acceleration in a non-relativistic shock (Beck & Krause 2005b)), L_ν is the luminosity of the source at frequency ν , V is the volume of the source and f is a filling factor describing the fraction of the volume occupied by emitting material.

Using the flux value measured at 608 MHz, a typical spectral index for a synchrotron source of $\alpha \approx -0.5$, a volume for the source of $V \approx 4 \times 10^{41} \text{ m}^3$ and a filling factor of $f = 0.5$, an equipartition magnetic field value of $B_{eq} \approx 690 \text{ } \mu\text{G}$ was obtained, larger than the previous value calculated in Ainsworth et al. (2014) of $110 \text{ } \mu\text{G}$.

A value for the equipartition magnetic field was also obtained based on the Razin effect fit. By equating the energy density in the nonthermal particles in the source U_{NT} with the energy density of the magnetic field U_B so that $U_{NT} = U_B$, an expression for B_{eq} dependent on the parameters for the Razin effect fit can be obtained:

$$U_{NT} \approx (1+k) \frac{\kappa}{p-2} E_{\min}^{-p+2} \quad (2.9)$$

where $E_{\min}$ is the electron energy corresponding to the minimum frequency observed in the spectrum, and:

$$U_B = \frac{B^2}{2\mu_0} \quad (2.10)$$

The parameters κ and $E_{\min}$ are both dependent on B . By substituting the expressions for these parameters and then rearranging $U_{\text{NT}} = U_B$, an expression for the equipartition magnetic field strength can be obtained:

$$B_{eq}^{7/2} = \frac{82\mu_0}{p-2} \frac{5.36 \times 10^{25}}{(1.253 \times 10^{37})^{(p-1)/2}} \frac{K}{a(p)\Omega l} \left(\frac{\nu_{\min}}{C} \right)^{(-p+2)/2} \quad (2.11)$$

Using the values for K and p from the Razin effect fit, an equipartition magnetic field strength value of $B_{eq} \approx 376 \mu\text{G}$ was calculated. again larger than the value from Ainsworth et al. (2014), although smaller than the value calculated with standard equipartition magnetic field formula in the previous paragraph. From this value, an estimate for the nonthermal electron density in the equipartition regime can also be obtained by again integrating over the relevant range of electron energies. This gives an equipartition nonthermal electron density of $n_e^{\text{rel}} \sim 2 \times 10^{-7} \text{ cm}^{-3}$.

Though the value for B from Equation 2.4 from the Razin effect turnover frequency is significantly smaller than the equipartition values calculated, this is not unreasonable given the large uncertainties associated with the equipartition method, particularly regarding the volume of the source. In addition, it is not certain that the source components should be close to equipartition.

2.3.5 Knot D

Unfortunately, knot D was not detected at 608 MHz with the GMRT or in the 2012 VLA observations. However, an approximate spectrum was plotted including VLA archival data from 2015 at 2.5 GHz and with the VLA observation at 6 GHz from 2016. It seems to tentatively show a low-frequency turnover in its spectrum (see Figure 2.2). A Razin effect turnover was fitted to the spectrum and appears to be a plausible explanation for the low-frequency turnover in this emission knot as well. However, detections of this knot at only 4 frequencies, which are separated in time, make it difficult to be certain and so no detailed calculations were carried out. The spectrum appears to have a lower turnover frequency than knot C. If this low-frequency turnover is also caused by the

Razin effect, this could indicate different magnetic field strengths or electron densities in the two emission knots. More sensitive observations at other frequencies are required for any detailed analysis.

2.4 Conclusions

In this work, DG Tau A was successfully detected at 152 MHz using LOFAR. This is only the second time that a YSO has been detected at such low frequencies and the first time that LOFAR has detected nonthermal emission from a YSO. A low-frequency turnover was found in the synchrotron spectrum of an emission knot of the jet, the first time such a turnover has been detected in the jet of a YSO. Considering the possible mechanisms for the turnover, synchrotron self-absorption (SSA) can almost certainly be dismissed as a possibility given the implausible magnetic field strength required. Alternatively, the Razin effect seems to provide the best explanation as it provides the most realistic physical parameters for the emission knot.

If the Razin effect is the correct mechanism, an estimate can be obtained for the magnetic field in the emission knot of $B \sim 20 \mu\text{G}$ at a projected distance of $\sim 3 \times 10^3$ au from the central source. This would be the first time that the magnetic field strength along a YSO jet has been measured based on the low-frequency turnover in a synchrotron spectrum. Though it should be noted that the magnetic field strength in the jet could differ from that in the emission knot due to effects such as amplification and shock compression. In the future, this method could be applied to other YSO jets to measure the magnetic field strength at locations along the jet, although it requires the presence of nonthermal emission and knowledge of the electron density.

Unfortunately, there are currently no other known low-mass YSOs associated with nonthermal emission, though it has been detected in several high-mass YSOs. However, recent improvements in the sensitivity of radio interferometers such as the VLA, and the existence of sensitive low-frequency interferometers such as LOFAR may change this. In addition, the next generation of ultra-sensitive interferometers, such as the Square Kilometre Array (SKA) and the Next Generation Very Large Array (ngVLA),

could allow very sensitive studies of nonthermal emission to be carried out. Detecting the low-frequency turnover of synchrotron radiation may therefore prove to be a valuable method of measuring the magnetic field and complement other methods such as polarization measurements.

Chapter 3

Detection of coherent low-frequency radio bursts from weak-line T Tauri stars

T Tauri stars (TTs) are low-mass ($< 2 M_{\odot}$) young stellar objects (YSOs) that have evolved to the state where they are optically visible. They are frequently associated with magnetic activity seen in the form of flaring and magnetospheric accretion bursts (Feigelson & Montmerle 1999). T Tauri stars can be divided into two categories: classical T Tauri stars (CTTs), which are surrounded by a circumstellar accretion disk, and weak-line T Tauri stars (WTTs), where most, if not all, of the disk has dissipated. Flares from these objects are thought to be caused by magnetic reconnection events, similar to the ones observed on the Sun but with energies up to 10^6 times larger (e.g. Fernández et al. 2004). In addition, while reconnection in the Sun and other main sequence stars occurs in magnetic fields anchored in the star, reconnection in YSOs can occur in magnetic fields connecting the star and the disk (Waterfall et al. 2019).

Flaring of WTTs has been observed at a variety of wavelengths, from the X-ray (e.g. Uzawa et al. 2011) to radio regimes (e.g. Forbrich et al. 2017). In the radio band, observations have normally been made at centimetre wavelengths (≥ 1 GHz), where any detected emission is generally attributed to incoherent gyro-synchrotron emission

Table 3.1: Primary characteristics of KPNO-Tau 14 and LkCa 4.

	KPNO-Tau 14	LkCa 4
Spectral Type	M6	K7
Age (Myrs)	1 ¹	2 ⁶ / 0.5 ⁷
Distance (pc)	151.9 ^{+5.1} _{-4.8} ²	129.4 ± 0.6 ²
Mass (M _☉)	0.1 ¹	0.79 ± 0.05 ⁶ / 0.15 - 0.30 ⁷
Radius (R _☉)	~ 1 ³	2.3 ⁷
Rotation Period	1.86 ± 0.01 ⁴	3.37 ± 0.01 ⁸
Surface Magnetic Field Strength (kG)	-	1.6 ⁶
Coronal Temperature (MK)	12.6 ⁵	-

References. (1) Luhman et al. (2003); (2) Bailer-Jones et al. (2018); (3) Derived from Baraffe et al. (2015) models; (4) Scholz et al. (2018); (5) Güdel et al. (2007); (6) Donati et al. (2014); (7) Gully-Santiago et al. (2017); (8) Grankin et al. (2008)

Notes. Due to the uncertainty in the age and mass of LkCa4, two values are given for these parameters.

from energetic electrons excited in magnetic reconnection events (Waterfall et al. 2019).

Among main sequence stars, radio bursts have primarily been observed from M dwarfs, which are known to be magnetically active (Villadsen & Hallinan 2019). At frequencies below or around 1 GHz, such bursts have been found to be coherent emission, generated by either plasma emission or electron-cyclotron maser (ECM) emission. In both cases, coherent radio emission is generally characterised by very high brightness temperatures (Dulk 1985).

For example, one well-known flare star, UV Ceti, was detected at 154 MHz using the Murchison Widefield Array (Lynch et al. 2017). This emission was determined to be coherent based on its high brightness temperature ($10^{13} - 10^{14}$ K) and, most likely, ECM emission due to the presence of elliptical polarisation, which is thought to only be generated in the x mode (Dulk et al. 1994). From theory, it is predicted that ECM emission is entirely polarised in the x mode, while plasma emission is polarised in the o mode (Dulk 1985).

In addition, Zic et al. (2019) detected several periodic, coherent bursts from UV Ceti. The period of the bursts is consistent with the rotation period of the star, and, based on the presence of high brightness temperatures and elliptical polarisation, the bursts were determined to be strongly beamed ECM emission originating from a low-

density cavity within the magnetosphere of UV Ceti. This illustrates that auroral emission is possible in main sequence M dwarfs, similar to that seen in brown dwarfs and Solar System planets (Pineda et al. 2017; Zarka 1998).

Recently, coherent radio emission was detected from the flare star CR Draconis by Callingham et al. (2021a) using the Low-Frequency Array (LOFAR; Van Haarlem et al. 2013). When observed for a total of 6.5 days over one year, three bright radio bursts were detected in addition to an almost steady radio flux. Due to the high brightness temperatures, it was determined that the emission must be coherent. Moreover, it shows high levels of left-handed circular polarisation at all epochs. The consistent handedness of the circular polarisation indicates that the emission originates from a region with the same magnetic polarity. Thus, it is unlikely to be due to plasma emission generated by flares since the polarity would be expected to flip due to flares occurring across the stellar disk. In addition, elliptical polarisation was also seen in one of the detected bursts. Based on these factors, it was determined that the emission must be ECM emission.

While in its infancy, the study of YSOs at low radio frequencies is nevertheless a growing field thanks to new facilities such as LOFAR. Now one can utilise not only single target observations but also surveys such as the LOFAR Two-metre Sky Survey (LoTSS), which allow steady and transient sources to be found (Shimwell et al. 2017). The detection of emission at low frequencies can give us important information on the presence of any high energy particles and magnetic fields in YSO outflows (Ainsworth et al. 2014), provide evidence for possible star-planet interactions (Vedantham et al. 2020; Kavanagh et al. 2021), and help evaluate the importance of coronal mass ejection (CME) events in dissipating proto-planetary disks (Schnepf et al. 2015).

With these ideas in mind, an initial study of the Taurus Molecular Cloud (TMC) was performed using the High-Band Array (HBA) of LOFAR centred on approximately 150 MHz. It should be noted that the HBA was used for this study to take advantage of its much higher sensitivity compared to the LOFAR Low-Band Array (LBA). Here I discuss two WTTSs in the TMC, KPNO-Tau 14 and LkCa 4, the primary

characteristics of which are listed in Table 3.1.

KPNO-Tau 14 is a very low-mass ($0.1M_{\odot}$) WTTS of spectral type M6 (Luhman et al. 2003). This puts the star very close to the brown dwarf upper limit ($< 0.08M_{\odot}$). Based on *Gaia* measurements (Gaia Collaboration et al. 2018), its distance is $151.9^{+5.1}_{-4.8}$ pc (Bailer-Jones et al. 2018). No significant level of accretion is detected in this YSO and there is no infrared excess in its spectrum (Muzerolle et al. 2005; Mohanty et al. 2005), suggesting that there is no dusty disk present. Mohanty et al. (2005) found evidence that KPNO-Tau 14 may be a spectroscopic binary, based on the measurement of a double peak in the cross-correlation profile of its $H\alpha$ line. However, a high-resolution imaging survey by Kraus et al. (2006) using the Hubble Space Telescope (HST) was unable to resolve the system, giving an upper limit for the projected binary separation of <4 au. Magnetic activity has been inferred from KPNO-Tau 14. More precisely, X-ray emission was detected from this star in the XMM-Newton extended survey of the Taurus Molecular Cloud (XEST) by Güdel et al. (2007). During these observations, an X-ray flare was found with a luminosity of 1.58×10^{30} erg s $^{-1}$ (corrected for extinction) and a duration of > 6000 s (Stelzer et al. 2007). For comparison, the median X-ray luminosity for TTSSs detected by Güdel et al. (2007) was 0.7×10^{30} erg s $^{-1}$.

LkCa 4 is another WTTS of spectral type K7. Based on *Gaia*, the distance to this star is 129.4 ± 0.6 pc (Bailer-Jones et al. 2018). Like KPNO-Tau 14, it has no infrared excess, indicating that its dusty disk has dissipated or at least that significant grain growth has occurred, thus reducing its infrared emission (Esplin et al. 2014). LkCa 4 is known to be very magnetically active. This is deduced from its high level of starspot coverage and intense flaring. In fact, Gully-Santiago et al. (2017) discovered, using model atmospheres, that its spectrum is best fitted with two temperature components: a hot photosphere of $T_{hot} \sim 4100$ K covering 20% of the surface and a cool component with $T_{cool} \sim 2700 - 3000$ K accounting for the remainder. This suggests a remarkable starspot coverage of 80%. Flaring activity was previously seen at radio frequencies by Dzib et al. (2015), who observed mildly circularly polarised emission at 4.5 GHz and

Table 3.2: Details of the LOFAR observations of KPNO-Tau 14.

LoTSS pointing	Date	Time	Duration (hrs)	ν (MHz)	$\Delta\nu$ (MHz)	σ_{rms} (μJy)
-	2013-11-30	19:15:00	8.0	152	74	210
P068+26	2019-06-28	05:11:00	8.0	144	48	130

Notes. The Date and Time columns refer to the time in UTC at the start of each observation. ν is the central frequency of the observation, and $\Delta\nu$ is the bandwidth. σ_{rms} is the root-mean-square noise in Stokes I at the position of KPNO-Tau 14.

Table 3.3: Details of the LOFAR observations of LkCa 4.

LoTSS pointing	Date	Time	Duration (hrs)	ν (MHz)	$\Delta\nu$ (MHz)	σ_{rms} (μJy)	Phase
P061+29	2019-06-28	05:11:00	8.0	144	48	230	0.00 – 0.10
P062+26	2019-07-06	04:52:15	8.0	144	48	240	0.37 – 0.47
P065+26	2019-07-07	04:59:59	8.0	144	48	280	0.67 – 0.77
P064+29	2019-08-10	02:11:00	8.0	144	48	150	0.71 – 0.81

Notes. The Date and Time columns refer to the time in UTC at the start of each observation. ν is the central frequency of the observation, and $\Delta\nu$ is the bandwidth. σ_{rms} is the root-mean-square noise in Stokes I at the position of LkCa 4. The rotation phase is calculated using the known rotation period of LkCa 4 of 3.37 d (Grankin et al. 2008) such that phase 0.0 corresponds to the beginning of the observation of LoTSS pointing P061+29 (2019-06-28 05:11:00).

7.5 GHz. The emission is highly variable in Stokes I over three epochs, with changes of $92.7 \pm 22.7\%$ at 4.5 GHz and $90.8 \pm 20.1\%$ at 7.5 GHz. The circular polarisation fractions measured are 9.6% at 4.5 GHz and 18.8% at 7.5 GHz. LkCa 4 was also detected in X-rays by Chandra with a measured luminosity corrected for extinction of $2.0 \times 10^{30} \text{ erg s}^{-1}$ (Yang et al. 2012).

The magnetic field of LkCa 4 was modelled by Donati et al. (2014), who found that it possesses a strong and mainly axisymmetric magnetic field with a poloidal component of $\approx 2 \text{ kG}$ and a toroidal component of $\approx 1 \text{ kG}$. There is a large degree of uncertainty regarding the mass and age of LkCa 4 due to its high level of starspot coverage. While previous estimates, using an effective temperature of 4000 K, had found mass and age estimates of $0.79 \pm 0.05 M_{\odot}$ and 2 Myr, Gully-Santiago et al. (2017) find an effective temperature of 3180 K based on their two-temperature component model. Comparing this to stellar evolutionary models (Baraffe et al. 2015) gives mass and age estimates of $0.15 - 0.30 M_{\odot}$ and $\sim 0.5 \text{ Myr}$, respectively. These evolutionary models assume the star to be unspotted, however, and thus there remain major uncertainties.

In this chapter, I report on the serendipitous detection of bright radio bursts at low frequencies from these two sources using LOFAR. In section 3.1, I describe how the observations were carried out and the data reduction process. In section 3.2, I show the resulting images and light curves of the observations that were obtained and describe the characteristics of the bursts that were detected. In section 3.3, I then discuss which of the possible emission mechanisms can explain the observations. In section 3.4, I discuss whether the ECM mechanism could also be observed from YSOs at higher radio frequencies. Finally, in section 3.5, I present the concluding remarks of this chapter.

3.1 Observations and data reduction

3.1.1 2013 LOFAR observations

Although the main objective of the observation carried out on 2013 November 30 - December 1 which was discussed in chapter 2 was to detect emission from the vicinity of DG Tau A, the full field was examined for other YSOs. A catalogue of all radio sources down to a 5σ signal-to-noise threshold was extracted using PyBDSF (Mohan & Rafferty 2015), which was then compared with those of known YSOs in XEST (Güdel et al. 2007), the Gould Belt GHz Very Large Array (VLA) Survey (Dzib et al. 2015), the Taurus Spitzer Survey (Rebull et al. 2010), and the Taurus WISE Survey (Esplin et al. 2014).

Out of a total of 115 YSOs within the field, radio emission was detected at the position of 2 other YSOs in addition to DG Tau A: KPNO-Tau 14 and V1320 Tau, with the noise level ranging from $100 \mu\text{Jy beam}^{-1}$ at the centre of the field to $400 \mu\text{Jy beam}^{-1}$ near the edge of the field. While the flux of V1320 Tau is steady, KPNO-Tau 14 shows strong flaring during our observing run. As this work focuses on radio bursts, I will only discuss KPNO-Tau 14, leaving V1320 Tau for future work. Details of the observation are given in Table 3.2.

Due to inaccuracies in LOFAR beam models (van Weeren et al. 2016; Shimwell

et al. 2019), the derived flux density scale is often not accurate. In chapter 2, this was accounted for by comparing the integrated flux densities of several compact bright sources in the field with their values in the TIFR GMRT Sky Survey (TGSS; Intema et al. 2017). It was found that the average ratio of TGSS flux densities to LOFAR flux densities was 0.83 ± 0.12 . However, there are known to be significant uncertainties with the flux density scale of TGSS. For this reason, and also to be consistent with the 2019 observations in the next section, the correction factor for the flux density scale was re-derived for the work in this chapter using the NRAO VLA Sky Survey (NVSS; Condon et al. 1998) and the 6C catalogue (Hales et al. 1988). This technique is described in more detail in Hardcastle et al. (2021). The flux-scale correction derived is 1.04 ± 0.21 and so the flux densities measured in the LOFAR image were corrected by this factor. Shimwell et al. (2019) found that there is still, conservatively, a 20% error in the flux density scale of LOFAR images after correction. Therefore, there is estimated to be an error of ± 0.21 in the correction factor.

3.1.2 2019 LOFAR observations

The LOFAR Two-Metre Sky Survey (LoTSS) is mapping the entire northern sky at 120-168 MHz with a target sensitivity of $\sim 100 \mu\text{Jy beam}^{-1}$ (Shimwell et al. 2017). As part of this survey, the Taurus and Perseus Molecular Clouds were imaged in a co-observing proposal (LOFAR Project code:LC12_015) in a series of observations during LOFAR Cycle 12, from June – November 2019. These regions each contain at least 414 and 369 YSOs, respectively (Esplin et al. 2014; Young et al. 2015). Details of these observations are given in Table 3.2 and Table 3.3. Work on this survey is currently ongoing and the full results will be published in a future paper.

Direction-independent calibration was carried out on each of the observations using the PREFACTOR pipeline. To correct for DDEs, the data were processed using the DDF Pipeline (Shimwell et al. 2019; Tasse et al. 2020), which uses the KillMS package (Smirnov & Tasse 2015) to solve for the DDEs and then uses DDFacet (Tasse et al.

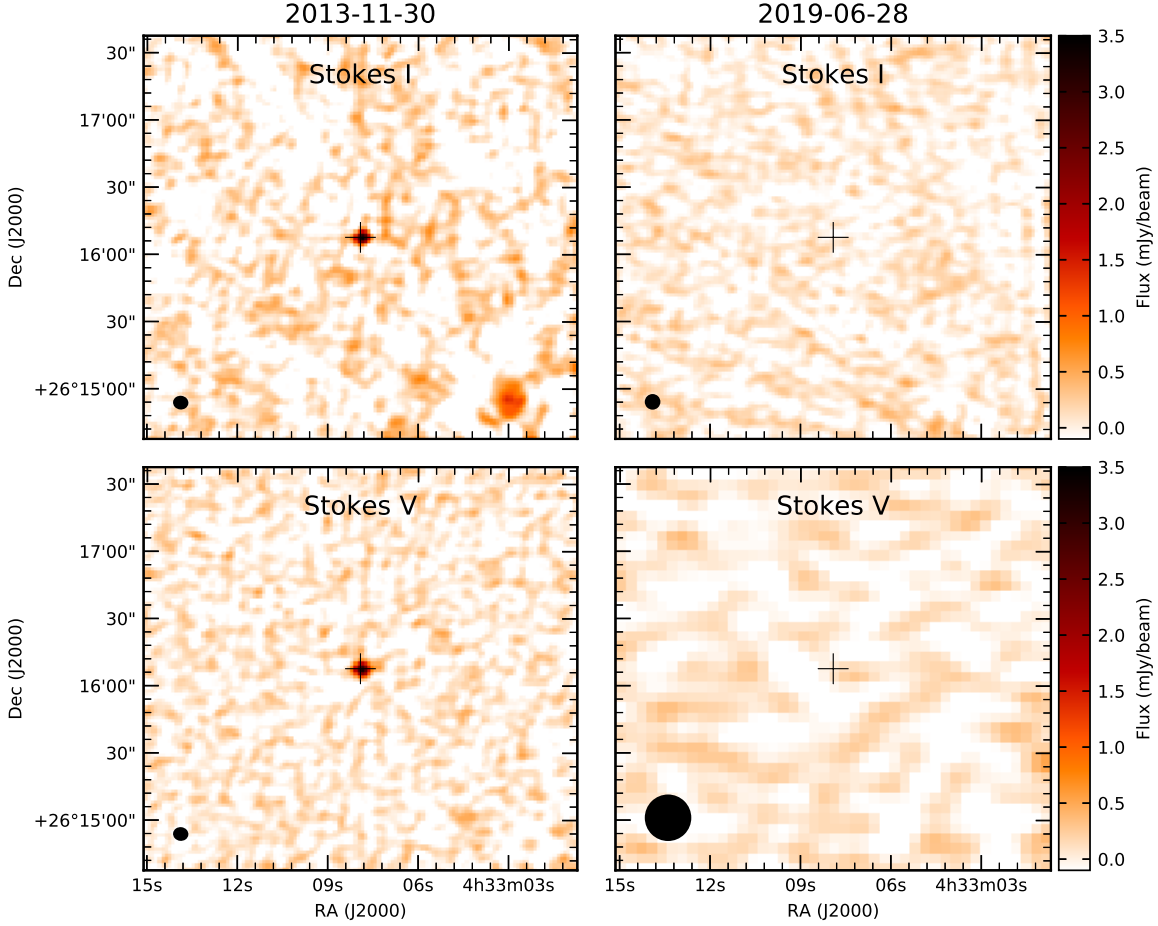


Figure 3.1: Stokes I (top) and Stokes V (bottom) images of KPNO-Tau 14 obtained from LOFAR at two epochs: November 30, 2013 (left) and June 28, 2019 (right). The restoring beam used for the 2013 images is $5.88'' \times 5.15''$. The restoring beams used for the 2019 LoTSS images are $6'' \times 6''$ and $20'' \times 20''$ for the Stokes I and Stokes V images, respectively. The optical position of KPNO-Tau 14 obtained from Gaia DR2 (Gaia Collaboration et al. 2018) is indicated by the cross.

2018), which applies the DDEs while imaging¹. The final output was a full-bandwidth Stokes I image of the field with a resolution of 6". In addition, a low-resolution 20" Stokes V image was also obtained.

As with the 2013 observations, and for each pointing, I extracted a list of sources for comparison with the previously cited YSO catalogues. In pointing P064+29, carried out on August 10, 2019, emission was detected towards LkCa 4. Similar to KPNO-Tau 14, this source also showed very strong flaring over the observing period. The position of LkCa 4 was also included in three other pointings: P061+29, P062+26, and P065+26, dated June 28, July 6, and July 7, 2019, respectively. However, no emission was detected towards LkCa 4 in any of these earlier epochs. Likewise, the position of KPNO-Tau 14 was included in the pointing P068+26, dated June 28, 2019, but no emission was detected.

Similar to the 2013 LOFAR observations, the flux density scales for each of the LoTSS pointings were corrected due to errors in the beam models by deriving correction factors using NVSS and the 6C catalogue. For P064+29, the flux-scale correction derived was 1.24 ± 0.25 and so the flux density measurements in this pointing were multiplied by this factor. Again, based on Shimwell et al. (2019) there was assumed to be an error of 20% in this correction factor.

3.2 Results

3.2.1 KPNO-Tau 14

In the 2013 observations, emission is detected from the position of KPNO-Tau 14 with a peak flux density (Stokes I) of 3.52 ± 0.73 mJy beam⁻¹, as seen in Figure 3.1. In addition, emission is also detected with a peak flux density of 3.28 ± 0.67 mJy beam⁻¹ in Stokes V, showing that the emission is highly circularly polarised with a polarisation fraction of $80^{+13}_{-17}\%$. Given that the source is unresolved, I take the peak flux density to be the total flux density.

¹See <https://github.com/saopicc> for the KillMS and DDFacet packages and <https://github.com/mhardcastle/ddf-pipeline> for the DDF-Pipeline

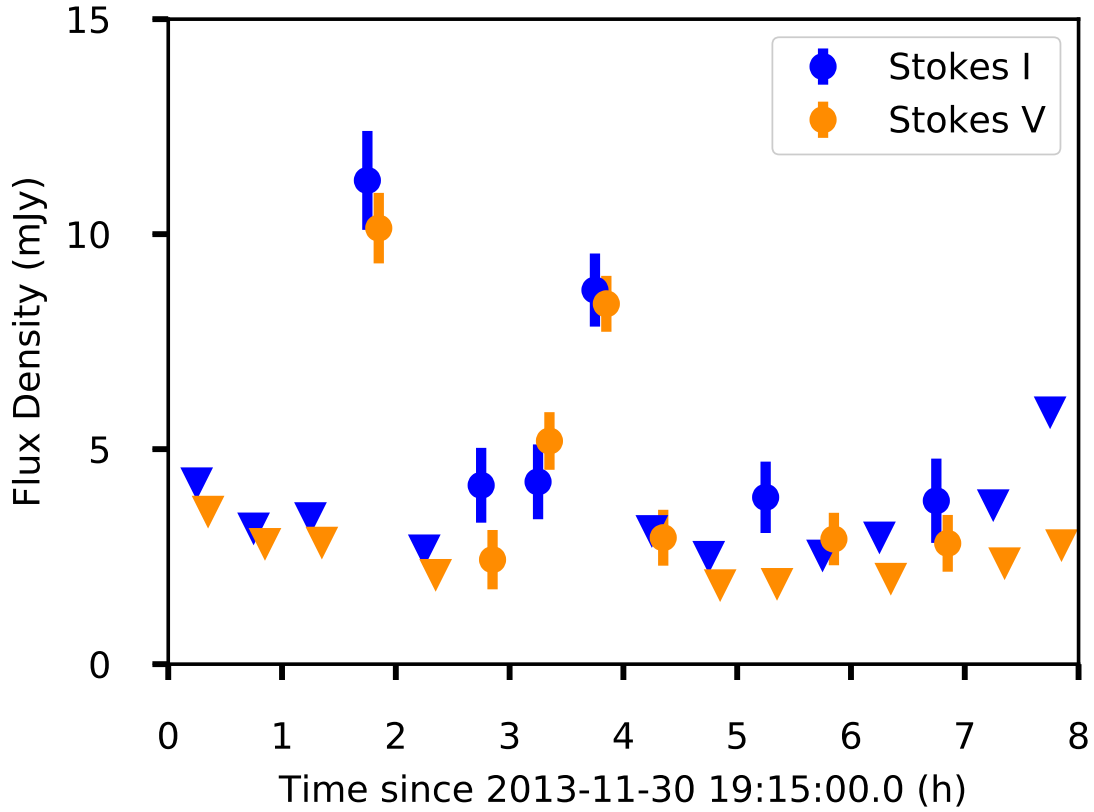


Figure 3.2: Light curve of KPNO-Tau-14 over the observation with a time resolution of 30 min. Stokes I and Stokes V measurements are shown by the blue and orange data points, respectively. The error bars are given by the root-mean-square noise σ_{rms} in the image. The Stokes V measurements are offset by +0.1 h with respect to the Stokes I measurements for better readability. Note that the absolute flux calibration error is not taken into account in the error bars as any absolute flux calibration offset should be the same for the whole observation. 3σ upper limits are shown by the triangles for the times in which no emission was detected from the source. See Appendix B for the flux density measurements and calculated polarisation fractions.

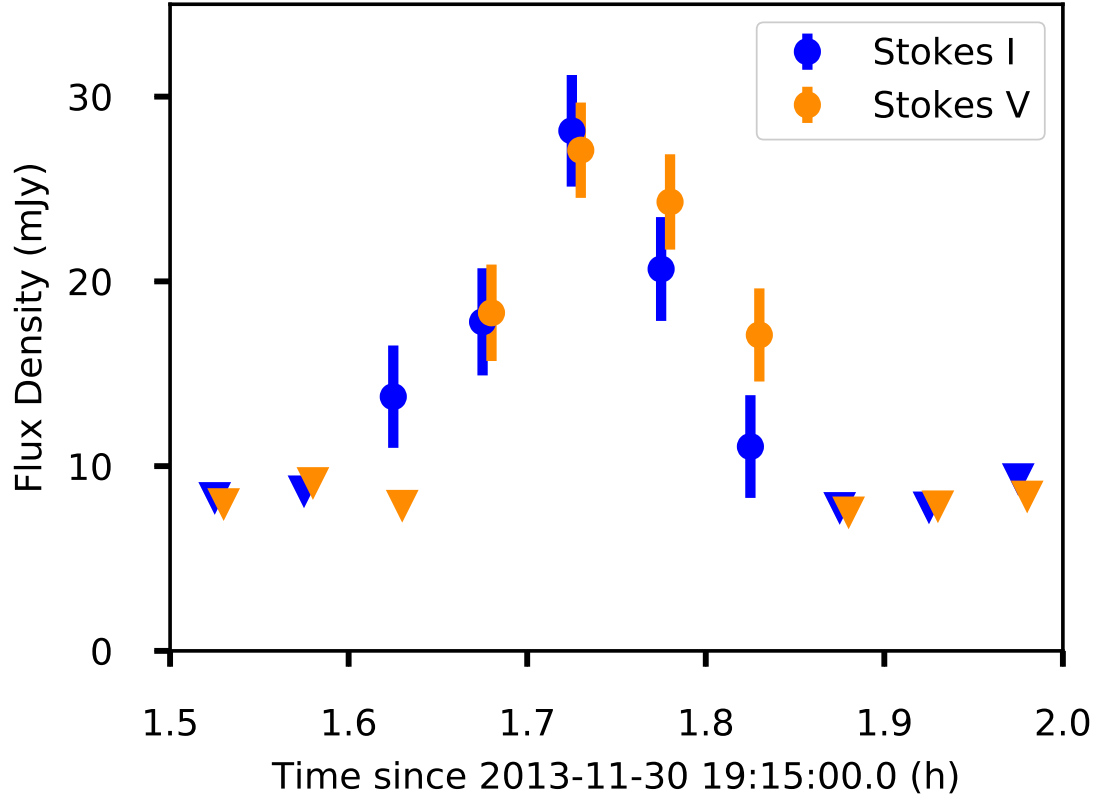


Figure 3.3: Light curve of the first burst detected from KPNO-Tau-14 with a time resolution of 3 min. Stokes I and Stokes V measurements are shown by the blue and orange data points, respectively. The error bars are given by the root-mean-square noise σ_{rms} in the image. The Stokes V measurements are offset by +0.005 h with respect to the Stokes I measurements for better readability. Note that the absolute flux calibration error is not taken into account in the error bars as any absolute flux calibration offset should be the same for the whole observation. 3σ upper limits are shown by the triangles for the times in which no emission was detected from the source. See Appendix B for the flux density measurements and calculated polarisation fractions.

The error in the flux density measurement, σ_{S_ν} , and subsequent flux density measurements in this chapter, was taken to be a combination of the root-mean-square noise σ_{rms} of the image and the absolute flux calibration error, found by Shimwell et al. (2019) to be 20% for LOFAR after applying a correction factor based on other surveys: $\sigma_{S_\nu} = \sqrt{\sigma_{rms}^2 + (0.20 \times S_\nu)^2}$. The polarisation fraction values and errors for this and all subsequent flux density measurements were found as follows. Gaussian distributions were generated for the Stokes I and Stokes V values, with a mean of the measured flux value S_ν and a standard deviation of σ_{S_ν} . The values in the Stokes V distribution were divided by the values in the Stokes I distribution to generate a distribution of polarisation fraction values. This polarisation fraction distribution was then truncated above 100% to remove any values higher than this. The value of the polarisation fraction was then taken to be the median of this truncated distribution and the errors give the 68% confidence interval.

To investigate time variability, the direction-dependent calibrated data were divided into time bins of 30 minutes. Each period was then imaged using WSClean in Stokes I and V. To check that the flux density scale calibration was accurate for all the images produced, the flux densities of several nearby compact bright sources in the field were measured in each image using PyBDSF and their light curves were plotted to check for any variation in the measured flux densities (see Appendix A for the light curves). Over the 8 hours, the flux densities of these sources vary with an average standard deviation of 12% relative to the mean flux density of each source. Given that this is within the typical absolute flux calibration error for LOFAR, I conclude that the flux density scale calibration is accurate for all of the images.

It was found that the flux density from KPNO-Tau 14 varied significantly over the 8 hours of observation. This can be seen in the light curve in Figure 3.2 where 2 bursts can be clearly seen. In particular, the first burst, 1.5 - 2 hours after the start of the observation, is very strong with a flux density of 11.25 ± 1.15 mJy. Note that the errors for this light curve and the subsequent light curves in this chapter do not take into account the absolute flux calibration error as any absolute flux calibration offset should

be the same for the whole observation.

In order to study the first burst in more detail, the data were further sub-divided into 3-minute time bins and plotted in the light curve seen in Figure 3.3, showing that the duration of the first burst is approximately 15 minutes and the flux density peaks at 28.15 ± 3.02 mJy. It should be noted that the true maximum flux density may be higher but due to the time binning, this may have been averaged out. Unfortunately, a higher time resolution cannot be obtained as the signal-to-noise would be too poor.

While the second burst seen in Figure 3.2 is not as strong as the first, it has a longer duration, lasting from 2.5 - 4.5 hrs after the start of the observation, and peaking at 8.70 ± 0.87 mJy. Some varying low-level emission is then detected during the rest of the observing window. The emission is highly circularly polarised throughout, with the exception of that detected at 5.0 - 5.5 hrs, and is consistently left-hand circularly polarised (see subsection 1.4.10). Both of the major bursts detected are strongly circularly polarised, with polarisation fractions of $88^{+8}_{-9}\%$ and $91^{+6}_{-8}\%$, respectively.

Based on the distance to the source, and assuming the emission is isotropic (i.e. not beamed), the first and largest burst has a peak luminosity of at least $(7.8 \pm 1.0) \times 10^{17}$ erg s⁻¹ Hz⁻¹, while the second smaller burst has a peak luminosity of at least $(2.4 \pm 0.3) \times 10^{17}$ erg s⁻¹ Hz⁻¹. To calculate the corresponding brightness temperatures, I assume, as indicative, the size of the emission region to be a circular disk with a radius equal to the radius of the star. From the Baraffe et al. (2015) evolutionary models for TTSSs, a $0.1M_{\odot}$ star with an age of 1 Myr should have a radius of $\sim 1R_{\odot}$. This would mean that the brightness temperatures of the first and second bursts are at least $(5.8 \pm 0.7) \times 10^{14}$ K and $(1.8 \pm 0.2) \times 10^{14}$ K, respectively. However, given that the emission region is likely much smaller than the entire stellar surface, the actual brightness temperatures are almost certainly being underestimated.

Finally, I remark that the presence of strong circular polarisation and the transient nature of the emission implies that the emission is associated with KPNO-Tau 14 and not a background extragalactic source for example. Also note that no emission is detected in either Stokes I or V from KPNO-Tau 14 in the later pointing (P068+26)

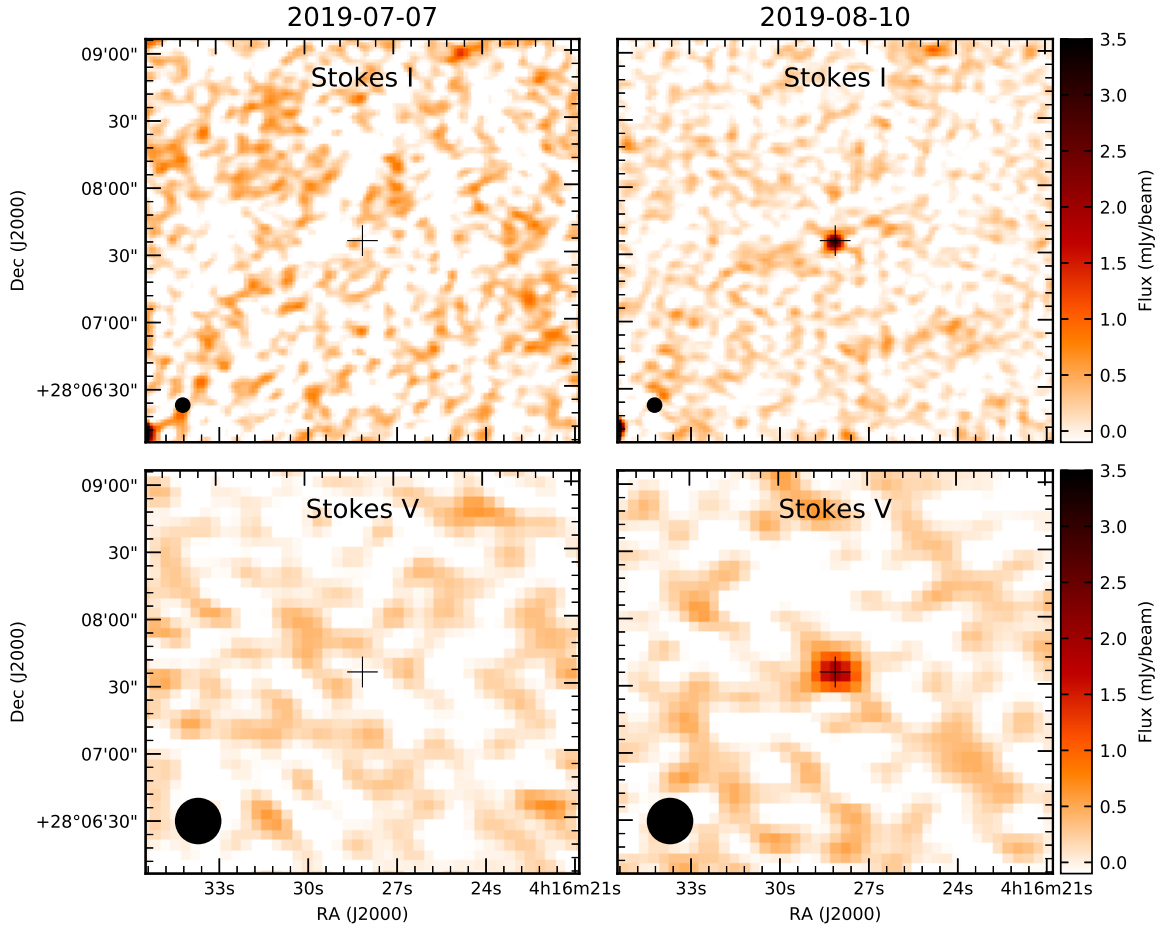


Figure 3.4: Stokes I (top) and Stokes V (bottom) images of LkCa 4 obtained from LOFAR at two epochs: July 7, 2019 (left) and August 10, 2019 (right). The restoring beams used for the images were $6'' \times 6''$ and $20'' \times 20''$ for the Stokes I and Stokes V images, respectively. The optical position of LkCa 4 obtained from Gaia DR2 (Gaia Collaboration et al. 2018) is indicated by the cross.

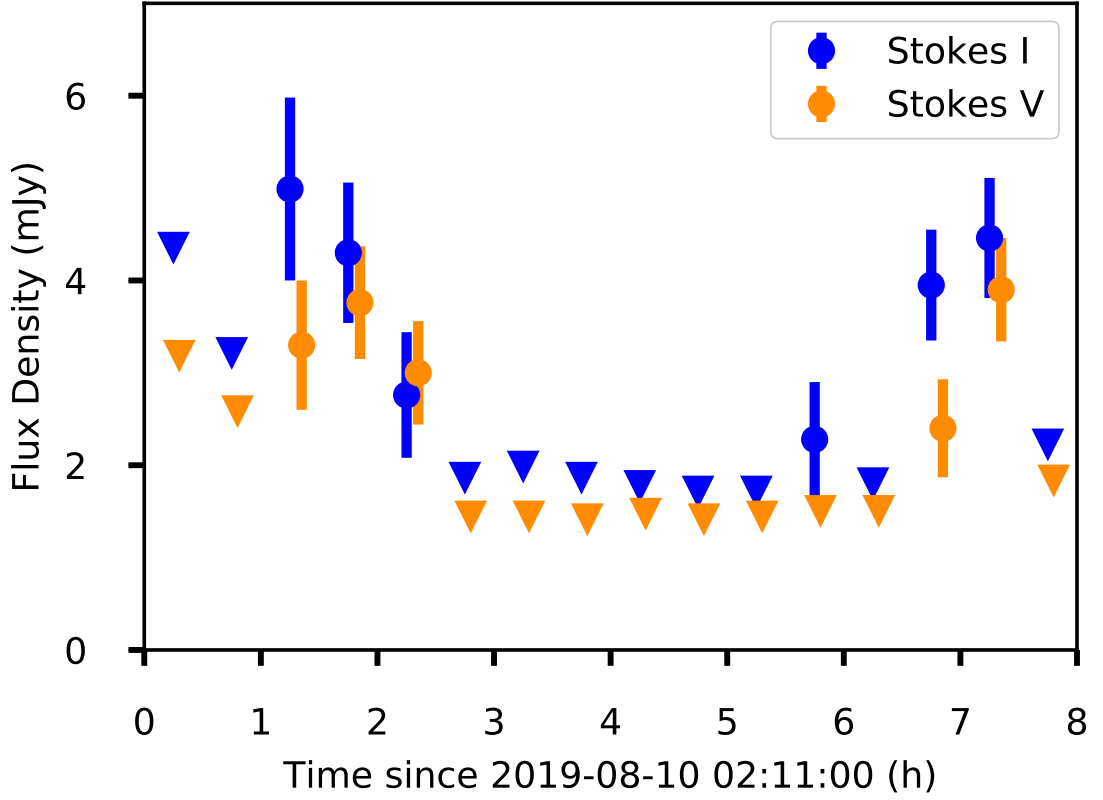


Figure 3.5: Light curve of LkCa 4 over the observation with a time resolution of 30 min. Stokes I and Stokes V measurements are shown by the blue and orange data points, respectively. The error bars are given by the root-mean-square noise σ_{rms} in the image. The Stokes V measurements are offset by +0.1 h with respect to the Stokes I measurements for better readability. Note that the absolute flux calibration error is not taken into account in the error bars as any absolute flux calibration offset should be the same for the whole observation. 3σ upper limits are shown by the triangles for the times in which no emission was detected from the source. See Appendix B for the flux density measurements and calculated polarisation fractions.

made in 2019, as seen in Figure 3.1, with a 3σ upper limit on the flux density in Stokes I of ≤ 0.39 mJy.

3.2.2 LkCa 4

During a LoTSS run on August 10, 2019, emission was found from LkCa 4 (pointing P064+29) with a peak flux density (Stokes I) of 3.12 ± 0.65 mJy beam $^{-1}$, as seen in Figure 3.4. In Stokes V, the corresponding value is 1.92 ± 0.42 mJy beam $^{-1}$, indicating a highly circularly polarised source with a polarisation fraction of $60^{+19}_{-15}\%$. Again, as

the source is unresolved, I take the peak flux density to be the total flux density.

As with KPNO-Tau 14, in order to detect changes in the emission, the direction-dependent calibrated data were divided into 30-minute time bins, each of which was then imaged using WSClean in Stokes I and V. The flux densities of several nearby compact bright sources in the field were measured in each image to check for that the flux scale calibration was accurate for all of the images used. Over the 8 hours, the flux densities of these sources vary with an average standard deviation of 8% relative to the mean flux densities of each source. Since this is well within the typical absolute flux calibration error for LOFAR, I conclude that the flux density scale calibration is accurate for all of the images (see Appendix A for the light curves).

As seen from the light curve in Figure 3.5, there are two bursts observed during the 8 hour run. The first lasts from 1.0 - 2.5 hours after the start of the observing period, peaking at 4.99 ± 0.99 mJy. The second lasts from approximately 6.5 - 7.5 hours, peaking at 4.46 ± 0.64 mJy. Both bursts are circularly polarised in a left-handed sense.

Based on the distance to the star, the peak luminosity of the first and second bursts are $\sim (1.0 \pm 0.2) \times 10^{17}$ erg s⁻¹ Hz⁻¹ and $\sim (0.9 \pm 0.1) \times 10^{17}$ erg s⁻¹ Hz⁻¹, respectively. Using a circular disk with a radius equal to the stellar radius $\sim 2.3 R_{\odot}$ (Gully-Santiago et al. 2017), as an indicative size of the source, the corresponding brightness temperatures are $\sim (1.6 \pm 0.3) \times 10^{13}$ K and $\sim (1.4 \pm 0.2) \times 10^{13}$ K, respectively. Similar to KPNO-Tau 14, however, it should be noted that the brightness temperatures are almost certainly much higher as the emission region is likely far smaller than the stellar disk.

Note that no emission is detected from LkCa 4 in the 3 other relevant LoTSS pointings, P061+29, P062+26, and P065+26, for which the corresponding 3σ Stokes I upper limits are ≤ 0.69 mJy, ≤ 0.72 mJy, and ≤ 0.84 mJy, respectively.

To check if the emission is related to the rotation period of the star, the rotation phases observed at each of the observations were compared. The phases were calculated based on the barycentric Julian dates (BJD) of the observations using the following

expression

$$E = \frac{BJD - 2458662.7}{3.37}, \quad (3.1)$$

where E is the rotational phase of the star, 3.37 ± 0.01 d is the rotational period of LkCa 4 (Grankin et al. 2008), and 2458662.7 is the BJD at the start of the first observation (P061+29), which I arbitrarily chose to correspond to a rotation phase of 0.0. The calculated rotation phases observed in each of the observations are listed in Table 3.3.

The two bursts observed in the LoTSS pointing P064+29 occur at phases $\approx 0.72 - 0.74$ and $\approx 0.79 - 0.80$, respectively. Therefore, the rotation phase corresponding to the first burst is observed in the observation of LoTSS pointing P065+26, which covers the phase range $0.67 - 0.77$. Note that there is an uncertainty in the phase of this observation relative to the observation of P064+29 of ± 0.03 due to the uncertainty in the rotation period of the star and the period of 1 month that elapsed between the two sets of observations. Given that the source is not seen in this observation it appears that the emission is not related to the rotation phase. Although it is important to note that the two sets of observations are a month apart and the bursting mechanism may be episodic.

In addition, detecting periodicity is also made difficult by potentially variable conditions in the source region. For example, for ECM emission variable maser conditions can lead to substantial variations in flux. In the case of CR Draconis, Callingham et al. (2021a) detect flux density variability of at least an order of magnitude, which is proposed to be due to ‘hotspots’ in the stellar magnetosphere, containing unusually dense and/or hot plasma, creating highly localised and variable emission. CM emission is generally also highly beamed, which may affect the ability to detect it. If the emission beam is frequently aligned away from the line-of-sight it will not be detected, making it difficult to identify any periodicity.

3.2.3 Frequency analysis

For the first burst from KPNO-Tau 14, the emission is sufficiently bright that a dynamic spectrum can be constructed. This was done by binning the data for the burst according to time and frequency and then imaging the region around KPNO-Tau 14 in Stokes I in each time-frequency bin and measuring the flux density of KPNO-Tau 14 in each image. The resulting dynamic spectrum is seen in Figure 3.6, with a time resolution of 6 min and a frequency resolution of 18.5 MHz.

Using this dynamic spectrum, the spectral properties of the burst can be analysed. It can be seen that the burst occurs over the entire bandwidth of the observation. The emission appears to be flat over the bandwidth, with a spectral index at 1.75 GHz of -0.03 ± 0.61 , and the flux density seems to rise and fall similarly across the bandwidth with no drift in the frequency of the emission apparent. However, it should be noted that the time and frequency resolution is very low, and so finer time-frequency structure cannot be ruled out. In addition, due to the low signal-to-noise ratio and non-detections in many of the time-frequency bins, significant changes in the spectrum of the burst could be hidden by the noise.

Unfortunately, for the other radio bursts detected, the flux density is not bright enough to obtain dynamic spectra due to the low signal-to-noise ratio.

3.2.4 Searching for linear polarisation

To search for linear polarisation, rotation measure (RM) synthesis (Brentjens & De Bruyn 2005) was used. This is required in order to correct for any Faraday rotation in the source, and/or from the foreground, an effect where the plane of linear polarisation rotates as it propagates through a magnetised plasma. The amount of rotation is frequency dependent:

$$\chi(\lambda) = \chi_0 + \phi\lambda^2, \quad (3.2)$$

where $\chi(\lambda)$ is the observed final angle, χ_0 is the initial angle, and ϕ is the Faraday depth, which depends on the average electron density and longitudinal magnetic field

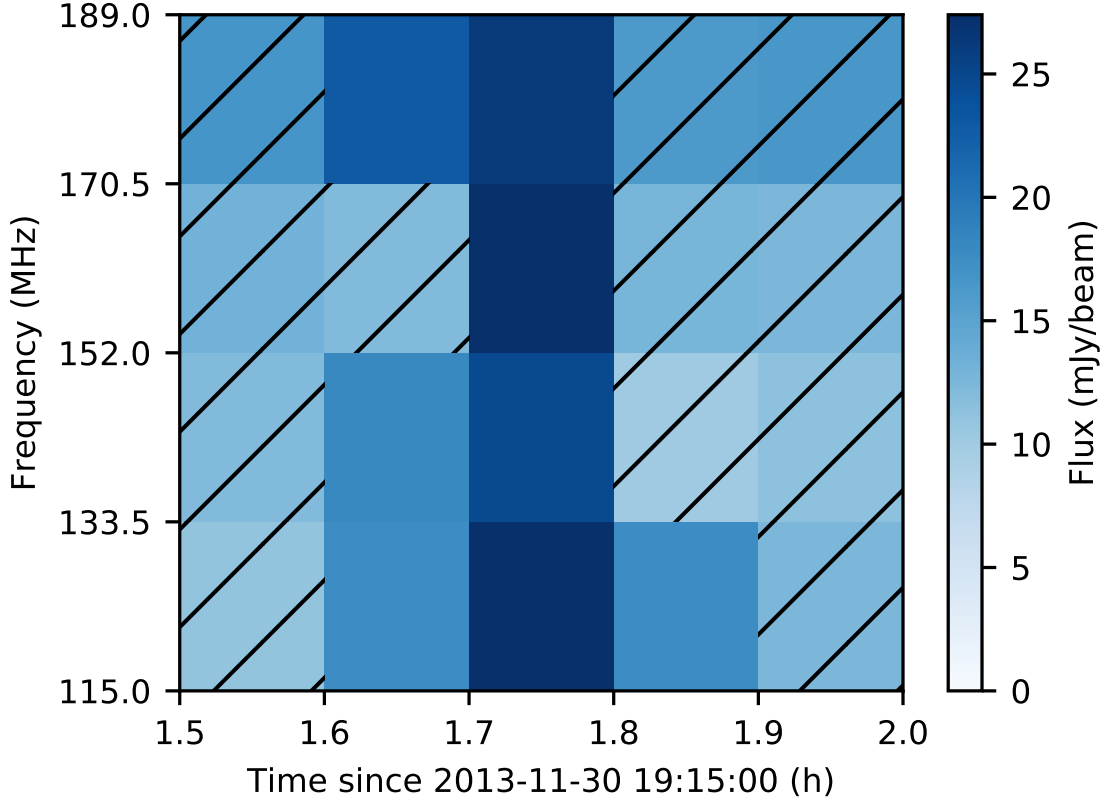


Figure 3.6: Dynamic spectrum of the first burst detected from KPNO-Tau 14. The burst was divided into time-frequency bins with a time resolution of 6 min and a frequency resolution of 18.5 MHz. The source was then imaged in each bin and its flux density was measured. 3σ upper limits are shown for the bins in which no emission was detected, indicated by the hatched lines. The mean noise of the images used for this spectrum is 4.4 mJy.

$\phi \propto \int n_e \mathbf{B} \cdot d\mathbf{r}$. As a result, when integrating a signal over a sufficiently wide bandwidth, the signal can become depolarised. This effect is particularly strong for LOFAR, as it operates at such long wavelengths and has a large fractional bandwidth. Rotation measure synthesis can get around this problem by imaging the Q and U images in narrow frequency channels and then performing a Fourier transform from the linearly polarised emission as a function of λ^2 to the linearly polarised emission as a function of ϕ .

When RM synthesis was done at the position of KPNO-Tau-14, no significant linearly polarised emission was found at any Faraday depth in the Faraday dispersion spectrum. The noise level in the Faraday dispersion is $189 \mu\text{Jy beam}^{-1} \text{RMSF}^{-1}$, where RMSF is the RM spread function, a point spread function in Faraday space due

to finite sampling in λ^2 -space. This gives a 3σ upper limit of $< 17\%$ on the linear polarisation fraction.

Similarly for LkCa 4, no evidence for linearly polarised emission was found. The noise level is $237 \mu\text{Jy beam}^{-1} \text{ RMSF}^{-1}$, giving a 3σ upper limit on the linear polarisation fraction of $< 35\%$.

However, this does not rule out the possibility of linear polarisation as it may not be possible to detect it due to lack of sensitivity. For example, Callingham et al. (2021a) detect linear polarisation from similar emission in CR Draconis with a linear polarisation fraction of $< 10\%$, which would be too faint to detect in the emission observed in KPNO-Tau 14 or LkCa 4.

3.3 Discussion

3.3.1 Determining the relevant emission mechanism

Both sources show a high level of circular polarisation in the bursts observed. In addition, their brightness temperatures are at least of the order of $\sim 10^{13} - 10^{15}$ K. This indicates a coherent emission mechanism, as incoherent mechanisms are limited to $10^{11} - 10^{12}$ K since above this, the source would be rapidly cooled by inverse Compton scattering (Kellermann & Pauliny-Toth 1969).

Assuming we are dealing with coherent radiation, the options are either plasma emission, as seen for example in the Sun from Type II radio bursts (Hariharan et al. 2014), or ECM emission, which is associated with auroral emitters such as planets, brown dwarfs and main sequence M dwarfs (Lynch et al. 2017; Pineda et al. 2017; Villadsen & Hallinan 2019; Zic et al. 2019; Vedantham et al. 2020; Callingham et al. 2021a) and possibly Type IV solar radio bursts (Treumann et al. 2011).

3.3.2 Plasma emission

Plasma emission occurs at the local plasma frequency, ν_p , (see Equation 1.22) and its second harmonic, but rarely if ever at higher harmonics. At the fundamental frequency,

plasma emission is expected to be polarised. However, the second harmonic is usually not polarised or only weakly polarised ($< 50\%$; Melrose et al. 1978). Given the high polarisation fraction of the emission in both of the sources, second harmonic plasma emission can therefore be ruled out.

The origin of plasma emission involves the production of longitudinal waves, known as Langmuir waves, in a plasma. These are usually generated by the injection of a hot plasma into a cooler ambient plasma. Some of the energy in the Langmuir waves is subsequently converted into electromagnetic waves.

Plasma emission, especially at higher frequencies, can suffer from free-free absorption due to overlying lower-density plasma blanketing the plasma in the source region (Dulk 1985). However, at low frequencies this is generally not an issue as the free-free optical depth is proportional to frequency: $\tau_{\text{ff}} \propto \nu^2$.

The maximum possible brightness temperature that can be generated by plasma emission can be estimated from the effective temperature of the Langmuir waves T_{L} , defined by (Dulk 1985):

$$W_{\text{L}} = \int \frac{k_{\text{B}} T_{\text{L}}}{(2\pi)^3} d^3 \mathbf{k}, \quad (3.3)$$

where W_{L} is the energy density of the Langmuir waves. It has been shown that the saturated value of W_{L} is 10^{-5} times the energy density of the background plasma (i.e. $W_{\text{L}} \approx 10^{-5} k_{\text{B}} T$; Benz 1993), where T is the temperature of the background plasma. From this, the limiting value of T_{L} can be shown to be (Dulk 1985):

$$T_{\text{L}} \lesssim 10^8 \frac{v_{\text{o}}^2}{c^2} \frac{v_{\text{o}}}{\omega_p} T, \quad (3.4)$$

where v_{o} is the velocity of the fast electrons in the hot plasma component and $\omega_p = 2\pi\nu_p$. Given that the brightness temperature of the plasma emission is limited to the effective temperature of the Langmuir waves, this also gives an upper limit on the brightness temperature.

From X-ray observations (Güdel et al. 2007), the average electron temperature in the corona of KPNO-Tau 14 is known to be 1.65×10^7 K. Using this as an estimate

for T and assuming a value of v_o similar to those seen in the Sun, $\approx 0.1c - 0.5c$, the brightness temperature is therefore limited to $T_b \lesssim 10^{15}$ K. While this is just about consistent with the estimated brightness temperatures of the bursts detected, it should be noted that those values were estimated based on an emission region equal to the size of the stellar disk and so are only lower limits. In reality, the emission region is likely much smaller and therefore the brightness temperature much higher.

This suggests that plasma emission is unlikely to be able to explain the bursts observed. Additionally, if the emission was coming from the entire stellar disk, this would imply the emission was due to flaring loops occurring across the entire stellar disk. If this were the case, we would not expect to obtain such high polarisation fractions as flaring loops from regions with oppositely directed magnetic fields would generate emission of oppositely handed polarisation, which would cancel each other out. Thus a high polarisation fraction such as that observed would require a smaller emission region.

Given that the temperature of the corona used is only an average, it is possible that the temperature could be higher in smaller regions of the corona. However, this runs into the issue again that the brightness temperature would then be required to be much higher in a smaller emission region and this would in turn require an even higher plasma temperature.

For LkCa 4, unfortunately no estimates of the ambient coronal temperature are available. However, assuming it is within the range typically seen for TTSs (i.e. $10^6 - 10^7$ K; Güdel et al. 2007), and taking the same estimates for v_o as before, the brightness temperature for plasma emission is limited to $T_b \lesssim 10^{14} - 10^{15}$ K. Again, this is marginally consistent with the estimated brightness temperatures of the bursts. Similar to KPNO-Tau 14 however, these estimated values are only lower limits based on an emission region equal in size to the stellar disk. As the real emission region is likely much smaller and therefore the brightness temperature likely much higher, it seems unlikely that plasma emission can explain the bursts.

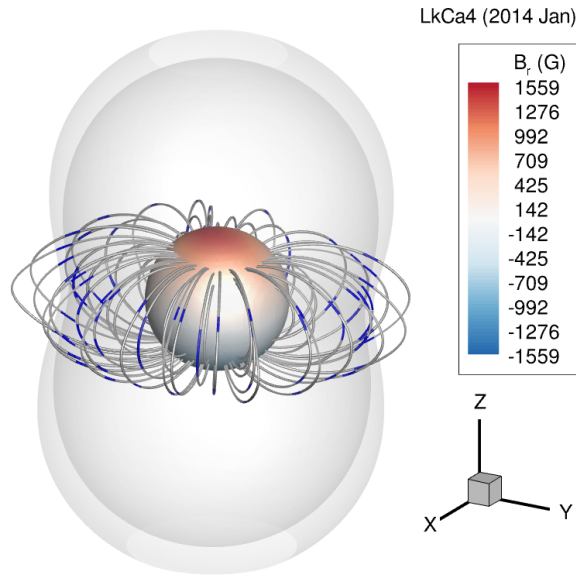


Figure 3.7: 3D magnetic map of LkCa 4, obtained by extrapolating the surface magnetic field maps obtained by Donati et al. (2014). The positive magnetic field strength values indicate where the magnetic field is pointing outwards from the surface, while the negative values indicate where it is pointing inwards towards the surface. The model used to calculate this magnetic map assumes that the field lines become fully radial at $3.4 R_*$. The grey surfaces indicate the region in the magnetosphere where the magnetic field strength is $40 - 70$ G. The locations along the magnetic field lines where the magnetic field strength is in this range are indicated by the blue line segments. Note that only the closed magnetic field lines are shown in this figure.

3.3.3 ECM emission

Electron-cyclotron maser emission is produced when radiation is amplified by stimulated emission at the local cyclotron frequency ν_c (see Equation 1.20 and its harmonics). In order for ECM emission to occur, the following conditions must be met: (a) a population inversion must exist in the electron energy distribution so that amplification by stimulated emission can occur, (b) the plasma frequency must be less than the cyclotron frequency, that is, either a strong magnetic field or a low-density plasma is required. Note that while the bandwidth of an individual pulse of ECM radiation is small, the frequency range of the emission as a whole can be quite large depending on the size of the emission regions and the range of magnetic field strengths that it spans.

One issue for ECM emission is the absorption of the emission by overlying plasma due to gyro-resonance absorption (Dulk 1985). This is particularly an issue for fundamental ECM emission as this is expected to be strongly absorbed at the second gyro-resonance layer (where $B = B_{\text{src}}/2$). One way to reduce this problem is to assume the emission is produced at the second harmonic. In this case, the emission only has to cross the third harmonic layer (where $B = 2B_{\text{src}}/3$). Alternatively, this issue can be avoided if the emission is generated in a low-density cavity within the magnetosphere, as this would give rise to a lower optical depth (Villadsen & Hallinan 2019; Zic et al. 2019).

If we are dealing with ECM emission, then the local magnetic field strength in the source region can be calculated. Given the observed LOFAR frequency range, and assuming the that it is fundamental ECM emission, the emission must be from a region with a field strength ranging from 40 G to 70 G.

For LkCa 4, Donati et al. (2014) show that the magnetosphere consists primarily of a 1.6 kG magnetic dipole component. Assuming a dipolar magnetosphere as an approximation, the distance of the emitting region from the star can then be estimated. This gives an emission region location at a height of $\approx 2R_*$. A more accurate estimate can be obtained by extrapolating the surface magnetic maps obtained by Donati et al. (2014) to obtain a 3-D map of the magnetosphere of LkCa 4, shown in Figure 3.7. The

grey surfaces indicate the region in the magnetosphere where the magnetic field strength is 40 – 70 G. The locations along the closed magnetic field lines where the magnetic field strength is in this range are indicated by the blue line segments, suggesting that the emission is from a region at a height of $\approx 2 R_*$.

No surface field strength estimates are available for KPNO-Tau 14. Assuming, however, that it is of the same order of magnitude as LkCa 4, then the emission should be from a similar height.

However, we are limited by the observation bandwidth. The frequency range of the ECM emission could extend well beyond that of LOFAR, which would imply larger emission regions and therefore a larger range of magnetic field strengths covered. For example, in the case of Earth and Jupiter, ECM emission has been observed with a frequency range of an order of magnitude or more (Melrose 2017).

Finally, it is worth repeating that for ECM emission to occur, the condition $\nu_p \lesssim \nu_c$ must be met (Dulk 1985). Therefore, an upper limit on the plasma density for both stars in their emitting regions can be obtained: $n_e \lesssim 10^8 \text{ cm}^{-3}$.

Coronal loops

One method of producing ECM emission is through a flaring coronal loop in the magnetosphere of the star. Heated thermal plasma is injected into the loop and creates a loss-cone distribution through magnetic mirroring at either end (Dulk 1985).

The brightness temperature generated by such a loop would depend on its size. To determine the dimensions of the loop needed to attain the observed brightness temperatures, I use the model given by Vedantham et al. (2020), which assumes a continuously operating loss-cone maser. This model suggests coronal loops of $10^3 R_*$ in the case of KPNO-Tau 14, and $10^2 R_*$ for LkCa 4. These values are far too large to be realistic suggesting the emission cannot be generated by this mechanism.

Co-rotation breakdown

Electron-cyclotron maser emission can be produced via the breakdown of co-rotation between the plasma surrounding the star and its magnetic field. This results in field-aligned currents that accelerate electrons into the corona and chromosphere of the star, creating the population inversion needed for ECM emission. Effectively this is a scaled-up version of the mechanism driving auroral radio emission in the case of Jupiter (Zarka 1998) and ultra-cool dwarfs (Nichols et al. 2012). More recently, this mechanism has been proposed to explain the LOFAR detected main-sequence M dwarf CR Draconis (Callingham et al. 2021a).

Both KPNO-Tau 14 and LkCa 4 are fast rotators, with rotation periods of 1.86 days (Scholz et al. 2018) and 3.37 days (Grankin et al. 2008), respectively. In addition, both show evidence of intense stellar activity, given their X-ray fluxes and, in the case of LkCa 4, high starspot coverage. Certainly, such activity could provide a source of plasma to drive the emission. In addition, WTTs are expected to have intense ionised winds with mass-loss rates, of the order of $10^{-12} - 10^{-9} M_{\odot}$ (Vidotto & Donati 2017), which is many orders of magnitude stronger than the Solar wind. Again such winds may prove to be a source of the plasma.

In the case of ultra-cool dwarfs and Jupiter, auroral radio emission is thought to occur at a wide range of heights along magnetic field lines from near the surface to several radii of the object (Nichols et al. 2012; Zarka 1998), giving rise to a wide frequency range of emission. If this is also the case for KPNO-Tau 14 and LkCa 4, this would be consistent with the inferred emission region at a height of $\approx 2R_*$ for LkCa 4. One would then also expect to see ECM emission at a wide range of frequencies, above and below that observed with LOFAR. For LkCa 4, with a measured polar magnetic field strength of 2 kG (Donati et al. 2014), this would imply ECM emission could, in principle, be detected at frequencies of up to 5.6 GHz (but see section 3.4).

Star-planet interaction

Another possible method for generating ECM emission is through the interaction of a stellar wind with the magnetosphere of an orbiting exoplanet. This is analogous to what happens in the Solar System with the giant planets (Farrell et al. 1999). For example Vidotto et al. (2010) and Lynch et al. (2018) predict that hot Jupiters orbiting young stars should produce the highest flux density of radio emission of any type of exoplanet.

From magnetohydrodynamical simulations of the stellar wind of the host star, Vidotto & Donati (2017) predict that the exoplanet V830 Tau b, a 2 Myr-old hot Jupiter orbiting a WTTS at 0.057 au, should produce an average radio flux density of 6 to 24 mJy, for a planet with a radius of 1 - 2 Jupiter radii. However, the flux density would not be constant, as it would vary over the planet's orbit as it moves through different regions of the stellar wind. This radio emission would be emitted at frequencies between 28 - 240 MHz, depending on the planetary magnetic field strength (10 - 100 G), meaning that such emission could potentially be detectable with LOFAR.

Given that V830 Tau is a WTTS, like KPNO-Tau 14 and LkCa 4, and is at a similar distance (147 pc), a hot Jupiter orbiting these sources could be able to generate similar levels of radio emission. In order for such emission to be observed using LOFAR at HBA frequencies, the surface magnetic field strength of the planet would need to be $\gtrsim 50$ G.

Alternatively, another method of generating radio emission through star-planet interaction is through a sub-Alfvénic interaction, similar to that seen in the Jupiter-Io interaction, which causes the Io-dependent decametric (Io-DAM) radio emission, in which the stellar magnetosphere couples with an orbiting body (Saur et al. 2013; Turnpenney et al. 2018). Magnetic energy flux is carried by Alfvén waves from the orbiting exoplanet and back towards the star, generating radio emission along the field line connecting the star and the planet. This was recently proposed to cause the radio emission seen with LOFAR in the nearby M dwarf GJ 1151 (Vedantham et al. 2020), and later confirmed through detection of a compatible exoplanet using the radial

velocity method (Mahadevan et al. 2021).

To check if this is a plausible explanation for the emission detected, I calculated the theoretical estimate for the power of the Alfvén waves generated by the interaction based on the following expression (Zarka 2007; Saur et al. 2013):

$$P_{\text{th}} = \frac{1}{2} R^2 v_{\text{rel}} (B \sin \theta)^2 M_A, \quad (3.5)$$

where R is the radius of the obstacle to the plasma flow (i.e. the planetary magnetosphere), v_{rel} is the relative velocity of the stellar wind with respect to the planet, B is the stellar magnetic field strength at the location of the planet, θ is the angle between the stellar magnetic field and v_{rel} , and M_A is the Alfvén Mach number at the position of the planet.

When calculating the theoretical power, the stellar wind is assumed to have the form of a Parker spiral (Parker 1958), which is given by the solution to the equation:

$$\frac{u_r(r)^2}{c_s^2} - \ln \left(\frac{u_r(r)^2}{c_s^2} \right) = 4 \ln \left(\frac{r}{r_c} \right) + 4 \left(\frac{r_c}{r} \right) - 3, \quad (3.6)$$

where $u_r(r)$ is the radial stellar wind velocity at the radial distance r from the surface of the star. The equation is parameterised by the sound speed:

$$c_s = \sqrt{\frac{k_B T}{m_p}}, \quad (3.7)$$

where k_B is the Boltzmann constant, T is the temperature of the stellar wind, and m_p is the proton mass. The wind passes through c_s at the critical distance r_c :

$$r_c = \frac{m_p G M_*}{4 k_B T}, \quad (3.8)$$

where G is the gravitational constant and M_* is the mass of the star.

The magnetic field components are given by:

$$B_r = B_0 \left(\frac{r_0}{r} \right)^2, \quad (3.9)$$

$$B_\varphi = B_r \frac{u_\varphi - \Omega r}{v_{\text{rel}}}, \quad (3.10)$$

where Ω_* is the stellar RV and u_φ is the azimuthal stellar wind velocity given by $u_\varphi(r) = (\Omega_* R_*^2)/r$.

The theoretically derived power was then compared with the observationally implied power based on the flux density observed:

$$P_{\text{obs}} = F \Omega d^2 \Delta\nu \epsilon, \quad (3.11)$$

where F is the observed radio flux density, Ω is the solid angle into which the emission is beamed, d is the distance to the source, $\Delta\nu$ is the total bandwidth of the emission, and ϵ is the efficiency of conversion of the energy of the vén waves to radio emission. The solid angle was estimated to be $\Omega = 1.6$ sr, similar to that of the Jovian radio emission (Zarka et al. 2004). The total bandwidth was taken to be the peak cyclotron frequency of the star $\Delta\nu \approx 2.8 B_0$ MHz, where B_0 is the surface magnetic field strength of the star in G.

For LkCa 4, again taking a surface magnetic field strength of 1.6 kG, this gives $P_{\text{obs}} \sim 6 \times 10^{27}$ erg s⁻¹ for the brightest burst observed. No magnetic field strength measurements are available for KPNO-Tau 14. Therefore, I used an estimate of 1 kG, which gives $P_{\text{obs}} \sim (2.8 \pm 0.4) \times 10^{28}$ erg s⁻¹ for the brightest burst observed. The efficiency of conversion was taken to be $\epsilon = 0.01$, similar to the value seen for the Io-DAM emission.

If, for example, there were a planet orbiting at $10 R_*$, it could be shown by using Equation 3.5 and Equation 3.11 that such values are appropriate to a planet with a magnetosphere of radius $0.6 R_J$, where R_J is the radius of Jupiter, for LkCa 4, or a planet with a magnetosphere of radius $2 R_J$ for KPNO-Tau 14. Such planets would likely be hot Jupiters. However, due to the many unknown parameters in this calculation, for example the solid angle of the emission, the efficiency of conversion, the orbital semi-major axis of the planet, these are only very rough estimates.

Based on these calculations, it seems that a star-planet interaction is at least a plausible explanation for the radio emission observed, either due to a stellar wind interaction or a sub-Alfvénic interaction. In order to determine whether or not the emission is of planetary origin, the YSOs would have to be monitored for a longer time to see if there is any periodicity in the radio emission unrelated to the rotation period of the star. A detection of an exoplanet around either source orbiting close to the star would also be a strong piece of evidence in favour of this. Unfortunately, detecting exoplanets in orbit around TTSs is difficult. Due to their high levels of activity, they generate large fluctuations in rotational velocity (RV). This makes it difficult to detect any orbiting exoplanets, even hot Jupiters.

Neither of the two sources is known to host any exoplanets. In the case of LkCa 4, Donati et al. (2014) attempted to detect a hot Jupiter by searching for periodic signals in its RV curve, after filtering out the RV variations due to the activity of the star. However, they were unable to find any signal with a semi-amplitude greater than 0.1 km s^{-1} , and conclude that there is no evidence for a hot Jupiter around LkCa 4, although they suggest that it is possible that there could be a planet with a similar orbital period to the rotation period of the star as its RV signal would then be hidden.

3.4 Detecting ECM emission in higher-frequency radio emission from YSOs

As stated earlier, until the availability of facilities such as LOFAR and the GMRT, most radio observations of YSOs have been made at centimetre wavelengths (Anglada et al. 2018). In these cases, what is observed on source tends to fall into two categories: either relatively steady (over at least several months) thermal bremsstrahlung emission arising from the ionised component of an outflow (Ray & Ferreira 2021a) or non-thermal flare-like emission, sometimes with a quasi-continuous background, that can vary over periods even less than an hour (Forbrich et al. 2017). The source of this emission is usually attributed to gyro-synchrotron radiation, produced by mildly relativistic

electrons with energies around a few MeV. While thermal emission from winds or outflows has only been seen from CTTSs (or their embedded counterparts) in the solar-mass regime, non-thermal emission is mainly observed from weak-line T Tauri systems, although it has also been seen in less evolved systems (Dzib et al. 2015). In all of these stars high energy processes are believed to be operating, however, in the case of less evolved systems, the non-thermal emission is expected to be absorbed by an overlying plasma (Andre 1996).

As ECM emission has been detected here at metre wavelengths from two YSOs, it is reasonable to ask whether some, or all, of the non-thermal radio flux in the centimetre band seen in other YSOs could be attributed to the same mechanism. The first point to note is that, if ECM emission is present, we expect to find a very high level of circular polarisation ($\approx 100\%$). In addition we would also anticipate very high surface brightness temperature measurements. Only a few detections of circular polarisation from YSOs at centimetre wavelengths exist and certainly some (Skinner 1993; Feigelson et al. 1998) appear consistent with optically thin gyro-synchrotron emission (Dulk 1985) where the circular polarisation fraction can reach at most a few tens of a per cent. That this is the appropriate mechanism is then backed up by the expected surface brightness temperature measurements. We are aware of just one YSO, T Tau Sb, where approximately 100% circular polarisation was observed on source at centimetre wavelengths (Smith et al. 2003) and another, EM* SR 20, in the decimetre regime (Pritchard et al. 2021). In at least the former case, the measured size of the emitting region, in combination with the flux, also suggests a coherent mechanism must be responsible. Further observations are clearly required to ascertain whether the type of emission observed here is common amongst YSOs at low frequencies or whether it is the rarity seen at higher frequencies.

3.5 Conclusions

I detect flaring radio emission from two WTTSs, KPNO-Tau 14 and LkCa 4, located in the TMC, at ~ 150 MHz using LOFAR. The emission detected from both sources

is strongly circularly polarised and highly variable over the course of each observation.

Based on the characteristics of the emission, in particular the high brightness temperature ($\gtrsim 10^{13} - 10^{15}$ K) and polarisation fraction (60 – 90%), I conclude that the emission must be due to a coherent emission mechanism, either plasma or ECM emission. Of the two coherent emission mechanisms, plasma emission seems unlikely as it cannot generate the high brightness temperatures and polarisation fractions observed, and so I conclude that ECM must be responsible.

In order to generate ECM emission, a breakdown of co-rotation between plasma surrounding the star and the stellar magnetosphere is the most likely explanation. Given that WTTSs are expected to have high mass-loss rates, this would provide a source of plasma to drive the emission. Such emission would be a scaled-up version of the auroral radio emission seen in Jupiter and ultra-cool dwarfs.

Alternatively, ECM emission could also be generated through a star-planet interaction. This could either be an interaction of the stellar wind with the planetary magnetosphere of an orbiting exoplanet, as seen in our own Solar System with the giant planets, or a sub-Alfvénic interaction of the stellar magnetosphere with the exoplanet, similar to the Jupiter-Io interaction and seen recently in the M dwarf GJ 1151. If a star-planet interaction is responsible, it could provide important information about exoplanetary magnetospheres. Additionally, given the difficulty in detecting exoplanets around YSOs using traditional methods, this could provide an alternative way of detecting young exoplanets.

These detections suggest that low-frequency radio observations are a valuable method for studying YSOs. As mentioned earlier, the Taurus and Perseus molecular clouds have recently been surveyed using LOFAR. Therefore, we may potentially detect similar emission from other YSOs in these clouds as part of this survey and other observations at low frequencies, giving us important information about their magnetic activity and potentially indicating the presence of exoplanets.

Chapter 4

A high-resolution radio study of the L1551 IRS 5 and L1551 NE jets

An important way of studying jets from YSOs is their free-free radio emission. Free-free (bremsstrahlung) emission at cm-wavelengths is commonly observed towards the base of the jet and is thought to trace material which has been ionized by internal shocks in the jet (Anglada et al. 2018). Radio observations at cm-wavelengths can therefore help to trace material which has recently been ejected by the YSO and calculate properties such as the direction, collimation, or ionized mass-loss rate of this recently ejected material. In addition, variations in the jet properties can be tracked on the timescale of only a few years. This can be compared with observations which track the outflow on multiple scales, such as molecular outflows or optical/IR jets.

In spite of the importance jets play in star formation, the mechanism by which these jets are launched and collimated is still somewhat uncertain (Ray & Ferreira 2021b). It is thought that material is accelerated by magnetocentrifugal forces along the magnetic field lines of either the star or disk. The two main models are known as the X-wind model (Shu et al. 1994) and the disk wind model (Blandford & Payne 1982). In the X-wind model, material is thought to be launched from close to the star at the co-rotation radius of the disk (~ 0.05 au) along the stellar magnetic field lines. In the disk wind model, material is launched from the surface of the disk at a range of

distances from the star up to several au along a magnetic field threading the disk. In both models, the material is then “self-collimated” by the helical magnetic field of the star/disk system.

In order to learn more about the launching and collimation of jets, high resolution observations near the base of the jet are very important. This makes radio observations of the free-free emission ideally suited for this purpose, particularly since optical or IR emission from this region is often obscured by high levels of extinction, particularly in more embedded sources. Previous observations at cm-wavelengths indicate that jets are collimated on scales of only a few tens of au or less (Anglada et al. 2018). The highest resolution observations of a low-mass protostar to date (Carrasco-González et al. 2019), suggest that in HL Tau the jet is already collimated at only ~ 1.5 au from the star.

In this chapter, I report on new high-resolution observations of two YSO systems in the L1551 star forming region: L1551 IRS 5 and L1551 NE. L1551 is a low-mass star forming region located in the south of the TMC at a distance of 147 ± 5 pc (Connelley & Reipurth 2018). The nearby location of this region makes it an ideal target for high-resolution observations as it allows us to resolve the regions close to the star. These observations were carried out using e-MERLIN and the VLA at a range of frequencies from 5 - 48 GHz. The resolution of these observations allows us to resolve emission on scales of only ~ 15 au.

L1551 IRS 5 is a well-known Class I source which has been classified as an FUor object based on spectral analysis (Connelley & Reipurth 2018). These are objects which show strong outbursts, in which their brightness increases by several orders of magnitude and are thought to be caused by rapid increases in the accretion rate of the star. Such outbursts can then last for decades (Audard et al. 2014). L1551 IRS 5 is a binary with a projected separation of $\sim 0.36''$ or ~ 50 au between the northern (N) and southern (S) component. The masses of the two components are $0.8M_{\odot}$ for the N component and $0.3M_{\odot}$ for the S component. Each has a circumstellar disk as well as a circumbinary disk surrounding the whole system (Cruz-Sáenz de Miera et al.

2019; Takakuwa et al. 2020). Both components have very high mass accretion rates of $\dot{M}_{\text{acc}} = 6 \times 10^{-6} \text{ M}_{\odot} \text{ yr}^{-1}$ and $\dot{M}_{\text{acc}} = 2 \times 10^{-6} \text{ M}_{\odot} \text{ yr}^{-1}$ for the N and S components respectively. Both of the components are known to drive jets which are seen in the form of thermal ionized jet emission in the radio (Rodríguez et al. 2003) as well as in the optical (Mundt & Fried 1983). The N jet has a position angle of $67 \pm 3^{\circ}$ while the S jet has a position angle of $55 \pm 1^{\circ}$ with the red-shifted portions of both jets pointing to the NE and the blue-shifted portions to the SW. L1551 IRS 5 was also the first source towards which a molecular outflow was detected (Snell et al. 1980). The inclination angle for both circumstellar disks is estimated to be $\sim 45^{\circ}$ and from this, we can assume the inclination angle for both jets to the line of sight is also approximately $i \approx 45^{\circ}$ (Lim et al. 2016b). Based on spectral line observations of the atomic outflow, the total mass loss rate of the system was estimated to be $\dot{M} = (1.3 \pm 0.8) \times 10^{-6} \text{ M}_{\odot} \text{ yr}^{-1}$ (Yang et al. 2022).

L1551 NE is another Class I binary system in the L1551 region located $\sim 2.5''$ northeast of L1551 IRS 5 comprised of a southeastern source (known as component A) and a northwestern source (known as component B) with an estimated projected separation of $\sim 0.5''$ or $\sim 70 \text{ au}$ between the two sources. The deprojected separation is calculated to be $\sim 145 \text{ au}$ (Takakuwa et al. 2014). The masses of the two components are estimated to be $\sim 0.68 \text{ M}_{\odot}$ for component A and $0.13 \text{ M}_{\odot}$ for component B. Similar to L1551 IRS 5, each component has a circumstellar disk as well as a circumbinary disk surrounding the whole system (Takakuwa et al. 2014, 2017). Thermal ionized jet emission is also seen from both components with position angles for the jets of $48 \pm 10^{\circ}$ and $58 \pm 8^{\circ}$ for components A and B respectively with the red-shifted portions of both jets pointing to the NE and the blue-shifted portions to the SW (Reipurth et al. 2002). The outflow from L1551 NE is thought to be responsible for the HH objects HH 28 and HH 29 and a series of HH knots known as HH 454 which surrounds L1551 NE. Lim et al. (2016a) calculated the inclination angles of both disks to $\sim 58^{\circ}$. From this, the inclination angle of the jet to the line of sight is estimated to also be $i \approx 58^{\circ}$.

The structure of this chapter is as follows. In section 4.1, I describe the radio

observations carried out using e-MERLIN and the VLA and the data reduction process. In section 4.2, I present the resulting images obtained and the flux densities measured at the different frequencies. In section 4.3, I model the spectrum of the sources and then use this to estimate the ionized mass-loss rate for each jet. I also discuss the variability seen in the jets compared to previous observations. Radiative transfer modelling is also performed for one of the jets in L1551 IRS 5. Finally, in section 4.4, I present the concluding remarks of this chapter.

4.1 Observations and Data Reduction

L1551 IRS 5 was observed using e-MERLIN on 2020 January 25 in C Band at 5 GHz/6 cm with a total bandwidth of 0.5 GHz (Project code: LE1007). The total time-on-source was ≈ 16 hrs. Using the VLA in the A configuration, the most extended configuration for the array, L1551 IRS 5 and L1551 NE were observed at C Band (6 GHz/5 cm), X Band (10 GHz/3cm), Ku Band (15 GHz/2 cm), and K Band (22 GHz/1.4 cm) on 2021 January 18, 15, 19, and 2020 December 29, respectively (Project code: 20B-122). All of the observations were processed and imaged using CASA. Self-calibration was also performed for the Ku band and K band observation. Unfortunately, self-calibration was not possible for the e-MERLIN observation or the VLA C band and X band observations due to insufficient signal-to-noise. As well as these observations we also used archival data for our sources from the Atacama Large Millimeter/submillimeter Array (ALMA). L1551 IRS 5 was observed in Band 3 (93 GHz/3.2 mm) and Band 4 (153 GHz/2 mm) on 2017 November 20 and 2017 November 04 (Project code: 2017.1.00388.S; PI: H. I. Liu), in Band 6 (225 GHz/1 mm) on 2017 July 24 (Project code: 2016.1.00209.S; PI: M. Takami), and in Band 7 (336 GHz, 0.9 mm) on 2017 July 27 (Project code: 2016.1.00138.S; PI: S. Takakuwa). L1551 NE was observed in Band 7 on 2019 October 07 (Project code: 2019.1.00847.S; PI: P. Sheehan). The ALMA images were obtained from the ALMA Science Archive. Details of all of the observations are given in Table 4.1 and Table 4.2.

In order to achieve good uv-coverage and create a high resolution image of L1551

Table 4.1: Summary of Observations of L1551 IRS 5

Instrument	Date	Time	λ (mm)	ν (GHz)	$\Delta\nu$ (GHz)	TOS (min)
e-MERLIN	2020-Jan-25	15:32:05	60	5	0.5	990
VLA	2021-Jan-18	06:34:54	50	6	4	6
VLA	2021-Jan-15	06:07:40	30	10	4	6
VLA	2021-Jan-09	00:38:00	20	15	6	80
VLA	2020-Dec-29	02:41:00	14	22	8	80
ALMA	2017-Nov-20	04:05:30	3.2	93	8	4
ALMA	2017-Nov-04	07:34:18	2.0	153	8	6
ALMA	2017-Jul-24	12:23:52	1.3	225	4	11
ALMA	2017-Jul-27	11:32:46	0.9	336	6	36

IRS 5 at C Band, the visibility data from the e-MERLIN C band observation and the VLA C band observation were combined. This takes advantage of the high resolution of e-MERLIN and the sensitivity and short baseline coverage of the VLA. As the observations are taken approximately a year apart, the proper motion of L1551 IRS 5 must be accounted for as well as any phase errors between the two observations. The two observations use different phase calibrators so there may be differences in the phase calibration and therefore in the positions of sources. The absolute proper motion of the system was calculated by Villa et al. (2017) based on VLA observations at 7 mm and found to be on the order of $\mu \sim 25 \text{ mas yr}^{-1}$ ($\mu_\alpha = +17.4 \pm 4.0 \text{ mas yr}^{-1}$, $\mu_\delta = -18.2 \pm 0.6 \text{ mas yr}^{-1}$). To account for any differences in position due to phase errors, the position of a point-like field source (RA: 04:31:43.00, Dec: +18:10:34.5) was compared between the two images. The source used was an extragalactic source and therefore should have no significant proper motion. The difference in the RA and Dec was found to be $\mu_\alpha = +20 \text{ mas}$ and $\mu_\delta = -70 \text{ mas}$ respectively. When combining the proper motion and the difference due to phase errors this gives a total difference in position between the two observations of $\mu_\alpha = +37.1 \text{ mas}$ and $\mu_\delta = -87.8 \text{ mas}$. Therefore, the e-MERLIN data was shifted so that the position of L1551 IRS 5 would match the position in the VLA data. Since the e-MERLIN and VLA observations have different bandwidths, only the part of the VLA bandwidth corresponding to the e-MERLIN bandwidth was used so that they would have the same central frequency and bandwidth when combined into one set of data.

Table 4.2: Summary of Observations of L1551 NE

Instrument	Date	Time	λ (mm)	ν (GHz)	$\Delta\nu$ (GHz)	TOS (min)
VLA	2021-Jan-18	06:34:54	50	6	4	6
VLA	2021-Jan-15	06:07:40	30	10	4	6
VLA	2021-Jan-09	00:38:00	20	15	6	80
VLA	2020-Dec-29	02:41:00	14	22	8	80
ALMA	2019-Oct-07	06:47:31	0.9	336	6	18

Table 4.3: Flux density measurements in L1551 IRS 5 N component and S component

ν (GHz)	L1551 IRS 5 N (mJy)	L1551 IRS 5 S (mJy)	σ_{rms} $\mu\text{Jy beam}^{-1}$	Synthesised Beam
5	0.11 ± 0.00	1.13 ± 0.10	11	$0''.22 \times 0''.11, -44^\circ$
10	0.51 ± 0.10	1.66 ± 0.07	10	$0''.20 \times 0''.18, 49^\circ$
13.5	0.90 ± 0.03	1.89 ± 0.04	6	$0''.17 \times 0''.13, -62^\circ$
17.5	1.19 ± 0.03	1.92 ± 0.03	6	$0''.14 \times 0''.11, -59^\circ$
20	1.27 ± 0.05	2.11 ± 0.06	12	$0''.12 \times 0''.09, -18^\circ$
24	1.63 ± 0.04	2.03 ± 0.05	12	$0''.094 \times 0''.074, -29^\circ$
93	34.1 ± 0.8	20.0 ± 0.5	50	$0''.124 \times 0''.089, 19^\circ$
153	85.1 ± 1.7	48.1 ± 0.8	76	$0''.058 \times 0''.041, -33^\circ$
225	271 ± 17	154 ± 15	750	$0''.15 \times 0''.13, -85^\circ$
336	497 ± 35	291 ± 27	1100	$0''.12 \times 0''.10, -65^\circ$

4.2 Results

4.2.1 L1551 IRS 5

L1551 IRS 5 was imaged and successfully detected in all bands. The flux densities were measured for the N and S components at each frequency using the task *imfit* in CASA and are listed in Table 4.3. Note that the Ku Band and K Band data were each split into two bandwidths given their large fractional bandwidths in order to obtain more data points. When measuring the total flux densities, the data at each frequency was reimaged, using the same outer uv-cutoff and synthesized beam for each band, to avoid any issues with differences in uv-coverage and synthesized beam size between the images. The uv-cutoff and synthesized beam used was the maximum baseline and the synthesized beam of the lowest resolution observation. For L1551 IRS 5, this was the X Band VLA observation which gives a uv-cutoff of $1.375 \times 10^6 \lambda$ and a beam of $0''.20 \times 0''.18$ with a PA of 49° .

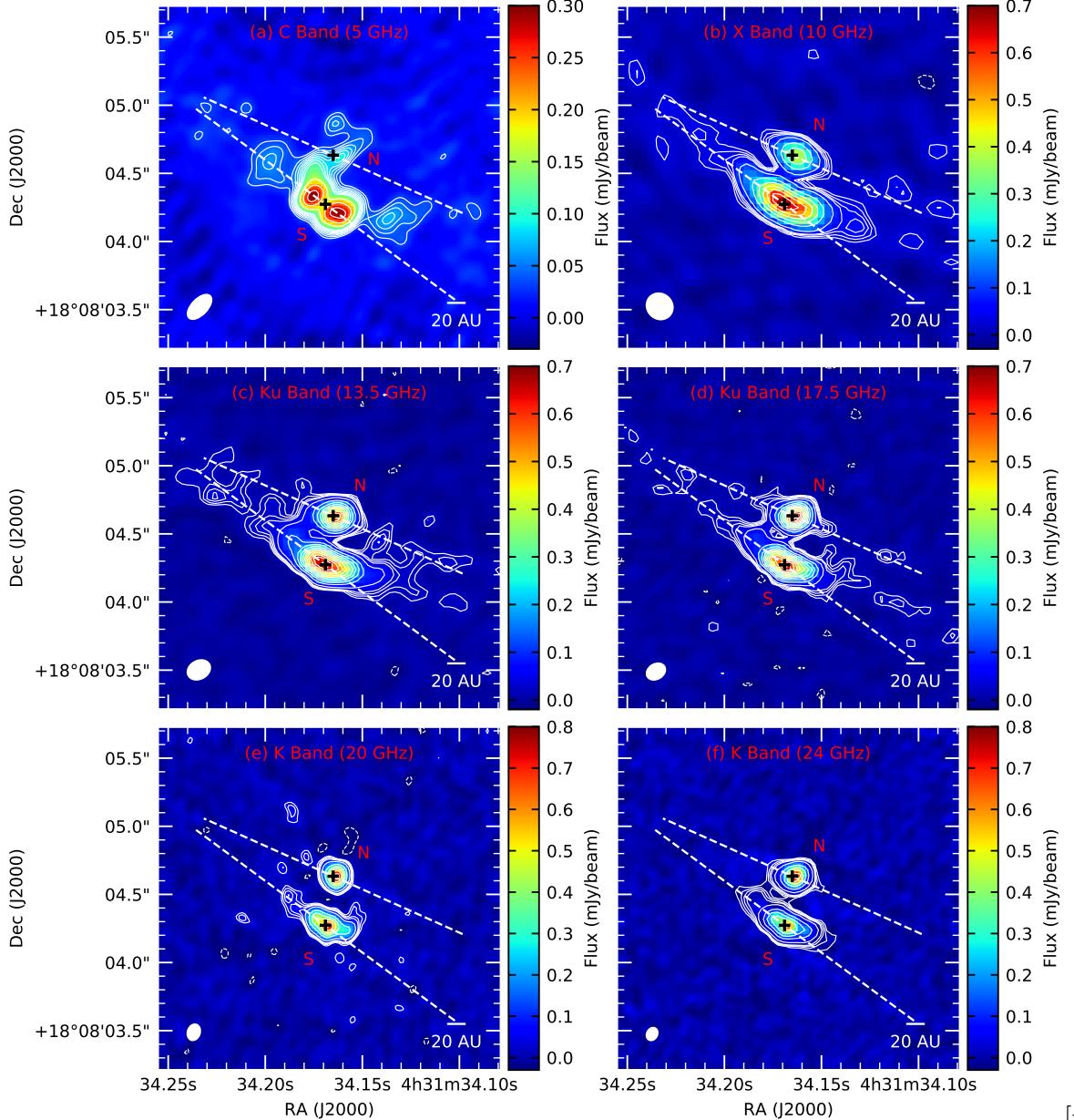


Figure 4.1: Images of L1551 IRS 5 from the VLA and e-MERLIN observations (contours + colourscale) as follows: (a) the e-MERLIN + VLA 5 GHz image (b) the VLA 10 GHz image (c) the VLA 13.5 GHz image (d) the VLA 17.5 GHz image (e) the VLA 20 GHz image (f) the VLA 24 GHz image. The black crosses in each image indicate the peaks of the disk emission of the two sources in the ALMA 153 GHz image to show the approximate positions of the central stars. The jet axes of the two components from Rodríguez et al. (2003) are shown by the dotted white lines. The contour levels in the 5 GHz image are $-3, 3, 4, 5, 6, 7, 8, 9, 10, 15, 20$, and $25 \times \sigma_{rms}$, where σ_{rms} is the root-mean-square noise of the image. For the 10 GHz image, the contour levels are $-3, 3, 4, 5, 6, 10, 20, 30, 40, 50$, and $60 \times \sigma_{rms}$. For the 13.5 GHz and 17.5 GHz images, the contour levels are $-3, 3, 4, 5, 6, 10, 20, 30, 40, 50, 60, 70, 80, 90, 100$, and $110 \times \sigma_{rms}$. For the 20 GHz and 24 GHz images, the contour levels are $-3, 3, 4, 5, 6, 10, 20, 30, 40, 50$, and $60 \times \sigma_{rms}$. The synthesized beam for each image is shown in the bottom left corner. The values for the synthesized beam and σ_{rms} of each image are given in Table 4.3.

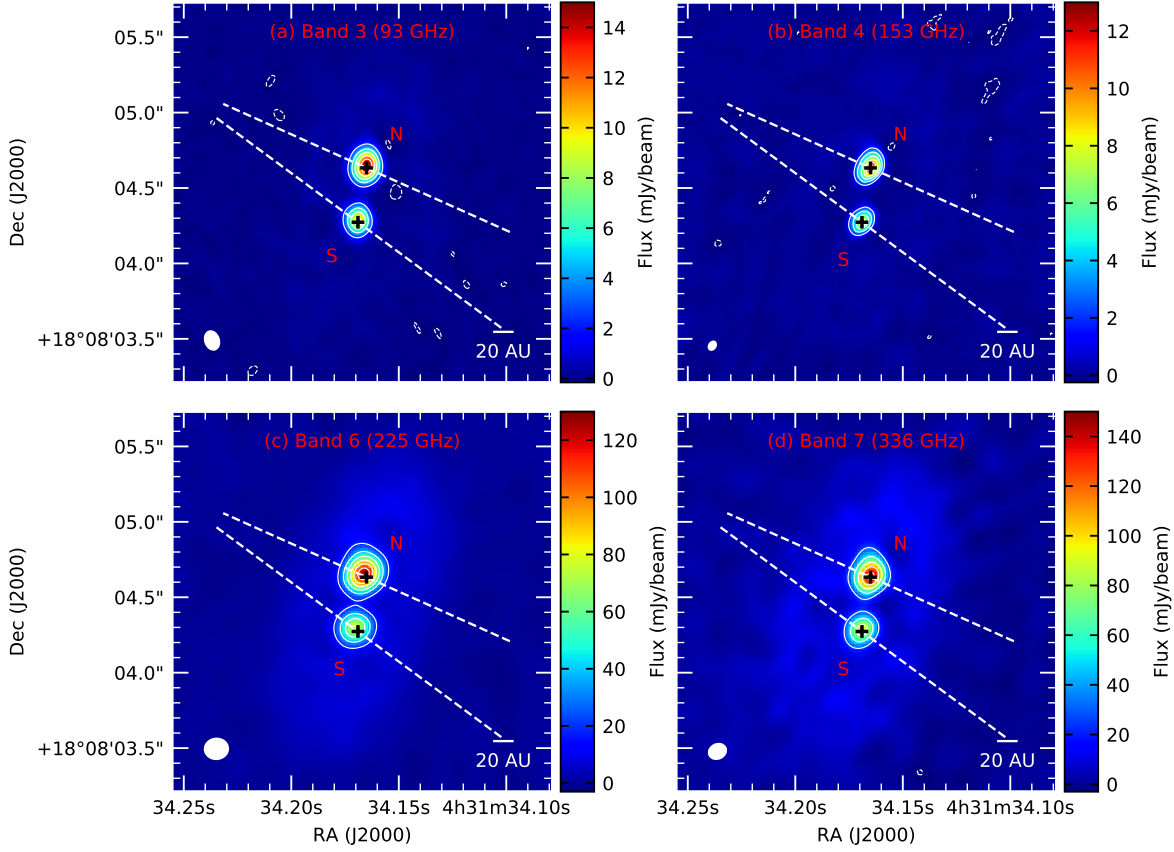


Figure 4.2: Images of L1551 IRS 5 from the ALMA observations (contours + colourscale) as follows: (a) the ALMA 93 GHz image (b) the ALMA 153 GHz image (c) the ALMA 225 GHz image (d) the ALMA 336 GHz image. The black crosses in each image indicate the peaks of the disk emission of the two sources in the ALMA 153 GHz image to show the approximate positions of the central stars. The jet axes of the two components from Rodríguez et al. (2003) are shown by the dotted white lines. The contour levels in the 93 GHz image are $-3, 50, 100, 150, 200, 250,$ and $300 \times \sigma_{rms}$, where σ_{rms} is the root-mean-square noise of the image. For the 153 GHz and 225 GHz images, the contour levels are $-3, 25, 50, 75, 100, 125,$ and $150 \times \sigma_{rms}$. For the 336 GHz image, the contour levels are $-3, 25, 50, 75, 100,$ and $125 \times \sigma_{rms}$. The synthesized beam for each image is shown in the bottom left corner. The values for the synthesized beam and σ_{rms} of each image are given in Table 4.3.

Table 4.4: Flux density measurements in L1551 NE component A and component B

ν (GHz)	L1551 NE A (mJy)	L1551 NE B (mJy)	σ_{rms} $\mu\text{Jy beam}^{-1}$	Synthesised Beam
6	0.32 ± 0.06	0.15 ± 0.03	16	$0''.37 \times 0''.30, 61^\circ$
10	0.36 ± 0.06	0.17 ± 0.05	26	$0''.20 \times 0''.18, 49^\circ$
13.5	0.60 ± 0.02	0.32 ± 0.02	6	$0''.17 \times 0''.13, -61^\circ$
17.5	0.83 ± 0.02	0.40 ± 0.02	6	$0''.14 \times 0''.11, -59^\circ$
20	1.16 ± 0.03	0.52 ± 0.03	11	$0''.10 \times 0''.09, -14^\circ$
24	1.42 ± 0.04	0.61 ± 0.04	11	$0''.084 \times 0''.073, -27^\circ$
336	372 ± 15	207 ± 22	1100	$0''.30 \times 0''.25, -12^\circ$

The images from the VLA observations are shown in Figure 4.1 in colourscale, with contours overlaid. The black crosses indicate the peak of the disk emission from the ALMA Band 4 image (153 GHz/2.0 mm) to show the approximate positions of the central stars. These positions have been shifted, relative to the ALMA image, to account for the proper motion between the ALMA observation and the VLA observations based on the proper motion measurements from Villa et al. (2017). The dashed white lines indicate the position angles of the jets from Rodríguez et al. (2003).

The combined e-MERLIN and VLA C Band (6 cm/5 GHz) is shown in Figure 4.1 (a). At this frequency, it is expected that there should be very little dust emission from the source and thermal free-free emission from the jet should be dominant. As can be seen in the figure, emission is clearly detected along the jet axes of both components, consistent with most of the emission being from the jet rather than dust emission from the disk. Although for the N component, only very weak emission is detected along the jet just to the SW of the central star. In the S component, two main emission lobes are seen either side of the central star, with NE lobe corresponding the red-shifted jet and the SW lobe corresponding to the blue-shifted jet.

At the higher frequencies observed with the VLA (10 - 24 GHz), both thermal free-free emission from the jet and dust emission from the disk should contribute to the flux. This is consistent with what is seen. The emission of both sources clearly peak at the positions of the centre of the disks, consistent with dust emission. However, the emission is extended along the direction of the jet axes of both sources suggesting that there is free-free emission from the jet present. At ALMA frequencies, the emission

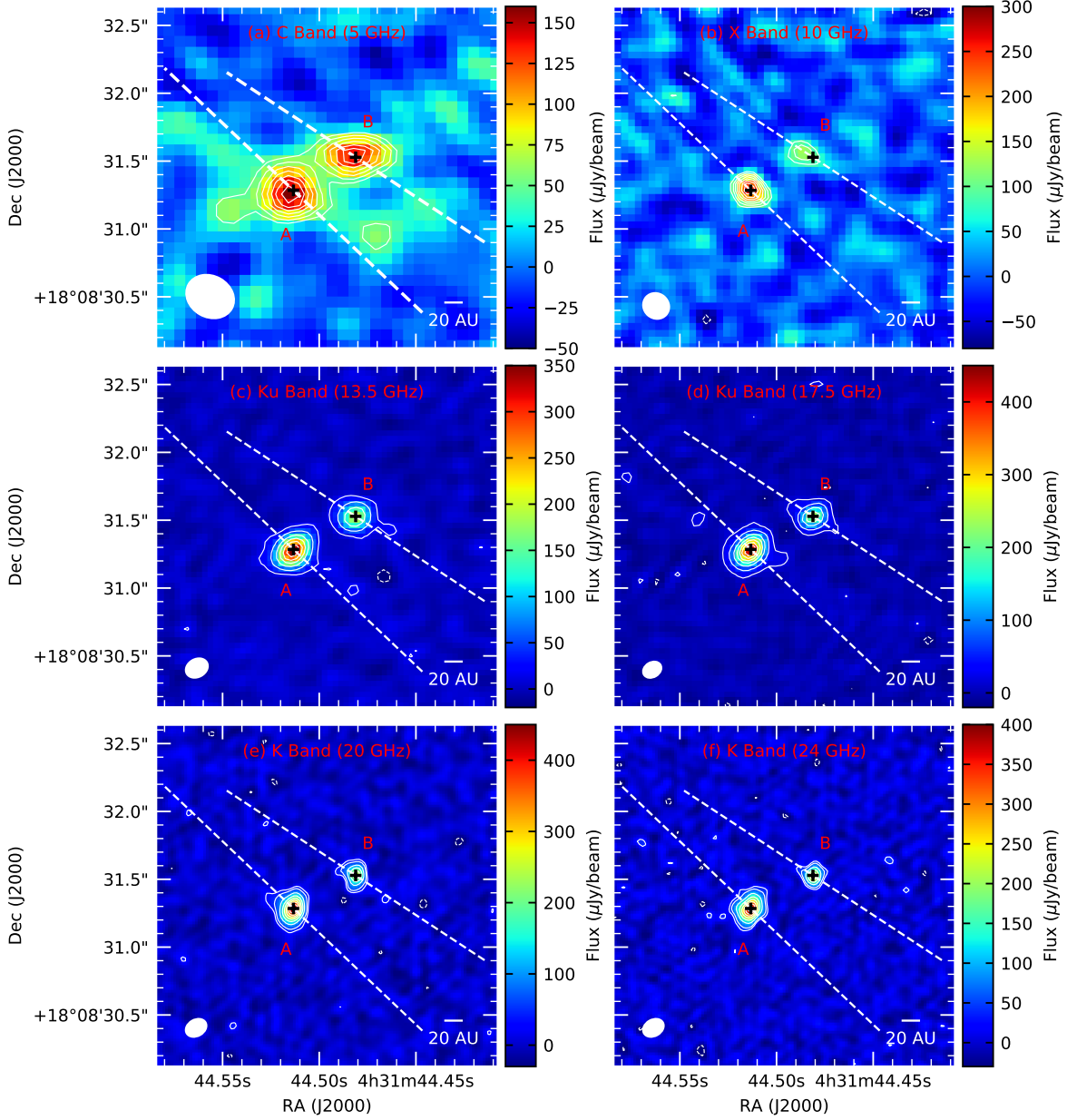


Figure 4.3: Images of L1551 NE from the VLA observations (contours + colourscale) as follows: (a) the VLA 6 GHz image. (b) the VLA 10 GHz image (c) the VLA 13.5 GHz image (d) the VLA 17.5 GHz image (e) the VLA 20 GHz image (f) the VLA 24 GHz image. The black crosses in each image indicate the peaks of the disk emission of the two sources in the 17.5 GHz image to show the approximate positions of the central stars. The jet axes of the two components from Reipurth et al. (2002) are shown by the dotted white lines. The contour levels in the 6 GHz image are $-3, 3, 4, 5, 6, 7, 8,$ and $9 \times \sigma_{rms}$, where σ_{rms} is the root-mean-square noise of the image. For the 10 GHz image, the contour levels are $-3, 3, 4, 5, 6, 7, 8, 9,$ and $10 \times \sigma_{rms}$. For the 13.5 GHz and 17.5 GHz images, the contour levels are $-3, 3, 10, 20, 30, 40, 50,$ and $60 \times \sigma_{rms}$. For the 20 GHz image, the contour levels are $-3, 3, 5, 10, 15, 20, 25, 30, 35,$ and $40 \times \sigma_{rms}$. For the 24 GHz image, the contour levels are $-3, 3, 5, 10, 15, 20, 25, 30,$ and $35 \times \sigma_{rms}$. The synthesized beam for each image is shown in the bottom left corner. The values for the synthesized beam and σ_{rms} of each image are given in Table 4.4.

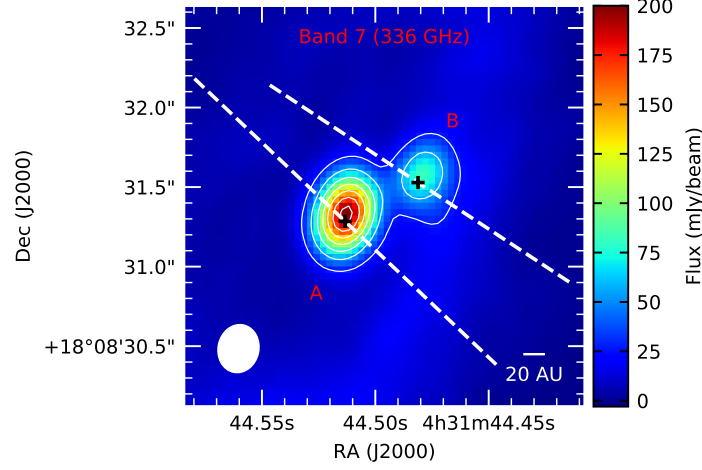


Figure 4.4: Image of L1551 NE from the 336 GHz ALMA observation (contours + colourscale). The black crosses in the image indicate the peaks of the disk emission of the two sources in the 17.5 GHz image to show the approximate positions of the central stars. The jet axes of the two components from Reipurth et al. (2002) are shown by the dotted white lines. Note that there is a displacement between the peaks of the ALMA 336 GHz image and the 17.5 GHz image due to proper motion of the source between the epochs (Oct 2019 and Jan 2021) which was not corrected for in the image. The contour levels are $-3, 25, 50, 75, 100, 125, 150, \text{ and } 175 \times \sigma_{rms}$, where σ_{rms} is the root-mean-square noise of the image. The synthesized beam for the image is shown in the bottom left corner. The values for the synthesized beam and σ_{rms} are given in Table 4.4.

is expected to be completely dominated by dust emission, with little or no free-free emission present. The ALMA images shown in Figure 4.2 agree with this, as the emission is centred at the positions of the stars at all frequencies, with no emission seen along the jet axes.

4.2.2 L1551 NE

L1551 NE was successfully imaged and detected in all of the bands in which it was observed. The flux densities measured for component A and component B at each frequency using the CASA task *imfit* and are shown in Table 4.4. Again, note that the Ku Band and K Band data were each split into two bandwidths in order to obtain more data points and when measuring the total flux densities, the data at each frequency was reimaged, using the uv-cutoff and synthesized beam corresponding to the maximum baseline and the synthesized beam of the lowest resolution observation. For L1551 NE,

this was the C Band VLA observation which gives a uv cut-off of $1 \times 10^6 \lambda$ and a beam of $0''.37 \times 0''.30$ with a PA of 61° .

The images from the VLA and ALMA observations are shown in Figure 4.3 and Figure 4.4 in colourscale, with contours overlaid. The black crosses indicate the peak of the emission from the Ku Band image (17.5 GHz/1.7 cm) to show the approximate positions of the central stars. The dashed white lines indicate the position angles of the jets from Reipurth et al. (2002). Similar to L1551 IRS 5, we should expect to see mainly free-free emission at lower frequencies, while at higher frequencies we should see a mixture of dust emission from the disk and thermal free-free emission from the jet. It should be noted that there is a displacement between the peaks of the ALMA Band 7 image and the 17.5 GHz image. This is likely due to proper motion of the source between the epochs (Oct 2019 and Jan 2021) which was not corrected for in the image as no proper motion measurements of L1551 NE were available.

4.3 Discussion

4.3.1 Spectrum

To learn more about the emission observed from these sources, the flux densities of the components of L1551 IRS 5 and L1551 NE were plotted as a function of frequency to give a spectrum for each source.

The resulting spectra for the N and S components of L1551 IRS 5 are shown in Figure 4.5 and Figure 4.6. For comparison, the flux density values given by Rodríguez et al. (1998) are shown in the figures as well. Both components were fitted with a model spectrum consisting of a combination of a thermal free-free emission component and a dust emission component. When modelling the free-free emission, the model from Reynolds (1986) was used. This assumes that the ionized jet begins at an injection radius r_0 , with an initial velocity v_0 , temperature T , width w_0 , and electron density n_0 . The width and electron density of the jet vary with distance along the jet as $w(r) = w_0(r/r_0)^\epsilon$ and $n(r) = n_0(r/r_0)^{q_n}$ respectively. The “standard” model for a

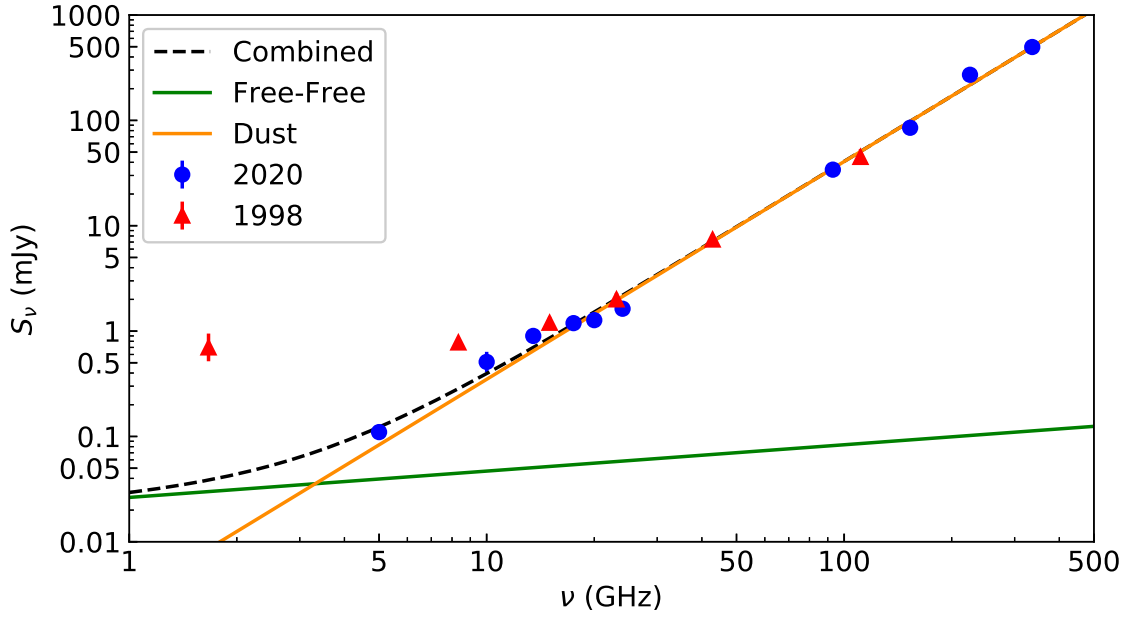


Figure 4.5: The spectrum of the N component of L1551 IRS 5 including the flux density values from this work obtained between 2017 - 2021 (blue circles) and the values from Rodríguez et al. (1998) (red triangles). The spectrum of the 2017 - 2021 data was fitted with a combination of free-free emission from the jet (green line) and dust emission from the disk (orange line). The combined fit is shown by the dashed black line.

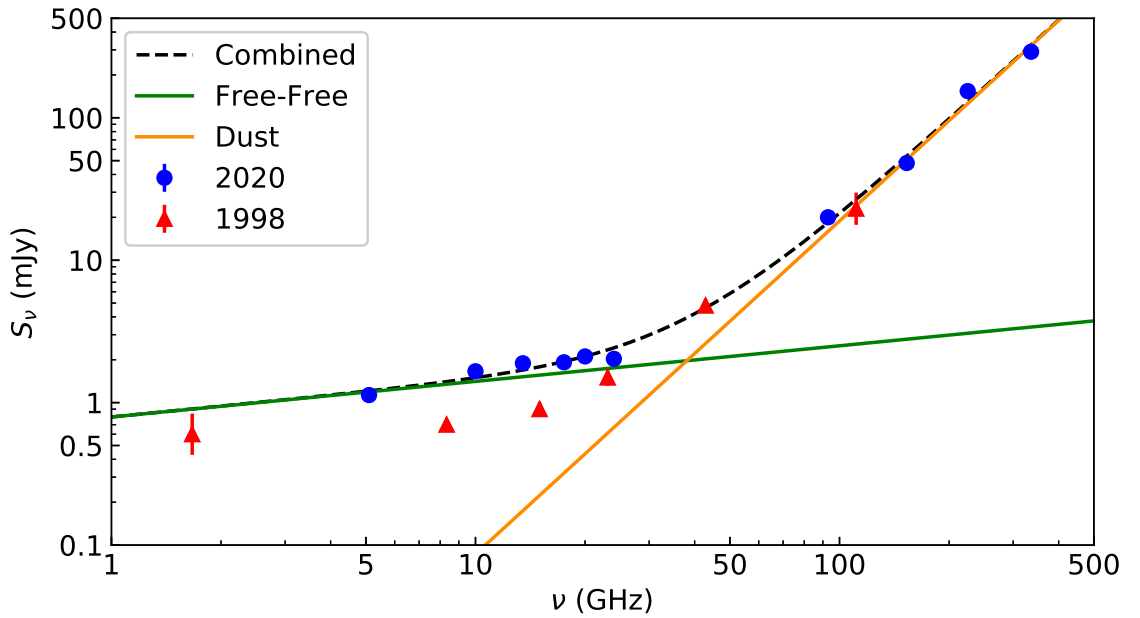


Figure 4.6: The spectrum of the S component of L1551 IRS 5 including the flux density values from this work obtained between 2017 - 2021 (blue circles) and the values from Rodríguez et al. (1998) (red triangles). The spectrum of the 2017 - 2021 data was fitted with a combination of free-free emission from the jet (green line) and dust emission from the disk (orange line). The combined fit is shown by the dashed black line.

collimated, ionized jet from Reynolds (1986) was used for which $\epsilon = \frac{2}{3}$, $q_n = -2\epsilon = -\frac{4}{3}$ and the temperature and velocity of the jet is constant ($T_e(r) = T_0$, $v(r) = v_0$). The spectrum model for the thermal free-free emission is then:

$$\begin{aligned} \left(\frac{S_\nu}{\text{mJy}} \right) &= 3.31 \times 10^{-9} \left(\frac{D}{\text{pc}} \right)^{-2} \left(\frac{T_e}{\text{K}} \right)^{-1/8} \left(\frac{r_0}{\text{au}} \right) \left(\frac{w_0}{\text{au}} \right)^{11/6} \left(\frac{n_0}{\text{cm}^{-3}} \right)^{5/3} \\ &\quad \times (\sin i)^{-1/6} \left(\frac{\nu}{\text{GHz}} \right)^{0.25} \\ &= K_{\text{ff}} \left(\frac{\nu}{\text{GHz}} \right)^{0.25} \end{aligned} \quad (4.1)$$

where D is the distance to the object and i is the inclination angle of the jet to the line of sight. When fitting the spectrum, all of the non-frequency parameters were combined into the parameter K_{ff} as a constant of proportionality for the free-free emission. Note that this equation is only accurate below the threshold frequency ν_{m} above which the jet becomes entirely optically thin. However, for typical parameters of a jet this is usually estimated to be $\nu_{\text{m}} > 40$ GHz (Anglada et al. 2018), at which point the spectra of the sources is clearly dominated by dust emission.

The dust emission component is modelled using simple power-law of $S_\nu \propto \nu^\alpha$ giving a total model spectrum of:

$$\left(\frac{S_\nu}{\text{mJy}} \right) = K_{\text{ff}} \left(\frac{\nu}{\text{GHz}} \right)^{0.25} + K_{\text{dust}} \left(\frac{\nu}{\text{GHz}} \right)^{\alpha_{\text{dust}}} \quad (4.2)$$

where K_{dust} is a constant of proportionality for the dust emission component and α_{dust} is the spectral index of the dust emission.

The S component is fit very well by this, as seen in Figure 4.6, with the flux density at low frequencies being mainly due to free-free emission and at higher frequencies being mainly due to dust emission with a spectral index for the dust emission of $\alpha_{\text{dust}} = 2.14 \pm 0.12$.

In the N component, the data is fit well by the model as seen in Figure 4.5. However, the fit seems to suggest that there is almost no free-free emission present and the emission is almost entirely due to dust emission with a spectral index of $\alpha_{\text{dust}} = 2.05 \pm 0.07$. Although it should be noted that in the Ku Band images of L1551 IRS 5 shown

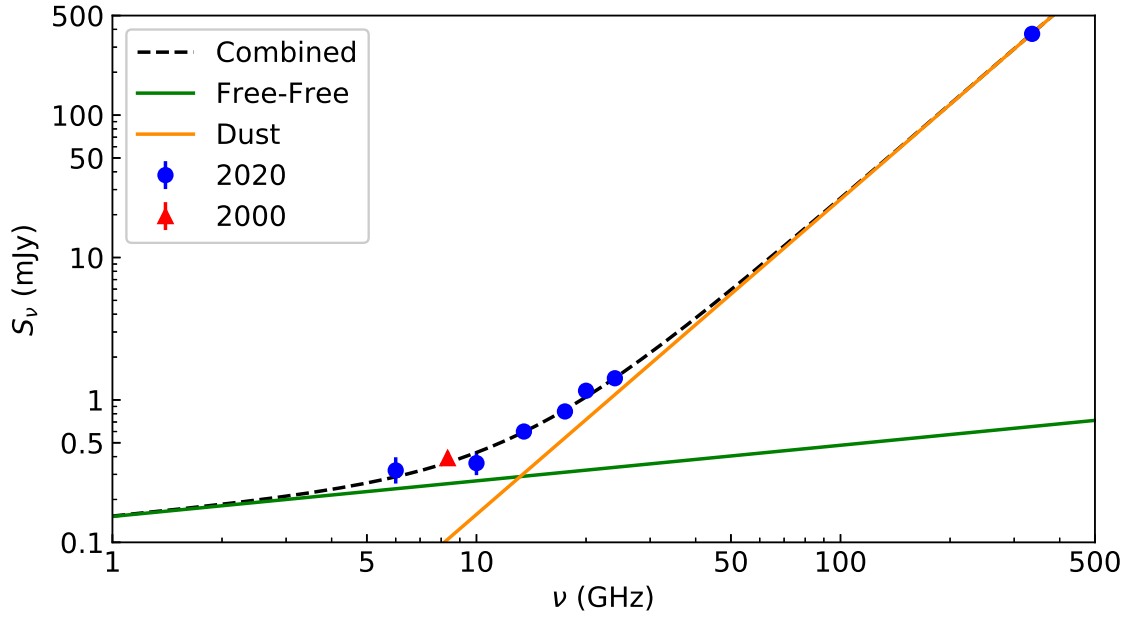


Figure 4.7: The spectrum of component A of L1551 NE. The spectrum was fitted with a combination of free-free emission from the jet (green line) and dust emission from the disk (orange line). The combined fit is shown by the dotted black line.

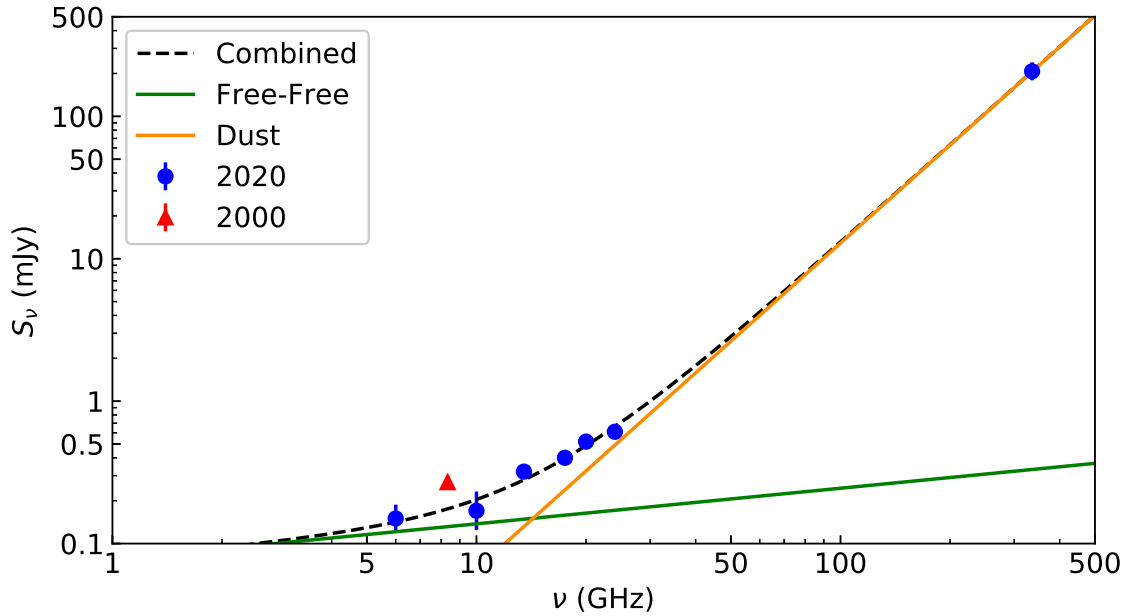


Figure 4.8: The spectrum of component B of L1551 NE. The spectrum was fitted with a combination of free-free emission from the jet (green line) and dust emission from the disk (orange line). The combined fit is shown by the dotted black line.

in Figure 4.1, there is clearly some emission seen along the jet axis of the N component, indicating that there is at least some level of free-free emission present.

For L1551 NE, unfortunately we do not have as many data points as L1551 IRS 5, particularly at higher frequencies where we only have ALMA observations in one band. However, it can still be seen in Figure 4.7 and Figure 4.8 that the spectra for both components A and B can be fit well by a combination of thermal free-free emission and dust emission, with free-free emission dominant at lower frequencies and dust emission dominant at higher frequencies. The spectral indices for the dust component are $\alpha_{\text{dust}} = 2.20 \pm 0.07$ and $\alpha_{\text{dust}} = 2.32 \pm 0.06$ for components A and B respectively.

4.3.2 Morphology

When comparing the image of L1551 IRS 5 obtained here with previous images at cm-wavelengths from Rodríguez et al. (1998) and Rodríguez et al. (2003), it is clear that there are significant changes in the morphology of the jets of both sources. For example, there are two strong emission knots seen along the jet axis of the N component in the image from Rodríguez et al. (2003), one to the NE and one to the SW. However, these seem to have disappeared in the images that we obtained. This is not that surprising when considering the proper motion of the material in the jet. If we assume a jet velocity of $v_{\text{jet}} \sim 200 \text{ km s}^{-1}$ for example, this would give a proper motion of $\sim 200 \text{ mas yr}^{-1}$ at a distance of 147 pc with an inclination of $i = 45^\circ$. Since the Rodríguez et al. (2003) observations were taken in February 2002, this would give $\sim 3.5''$ of proper motion between the two epochs.

One thing of interest to note, is that in Rodríguez et al. (2003), they detect a region of emission between the two jets $\sim 0.6''$ to the SW of the two sources. They speculate that this emission could be due to a region of shock interaction between the outer parts of the outflows. They state that to test this hypothesis, a lack of proper motion would have to be observed in this emission as an interaction zone between the two outflows should remain stationary. In the 5 GHz image that we obtained, a region of emission is also detected between the outflows to the SW of the two sources at a similar position to

that of the emission detected by Rodríguez et al. (2003). If this emission corresponded to the same emission region as the one they detected, this would indicate a lack of proper motion and back up their hypothesis of this being a region of shock interaction between the two outflows.

4.3.3 Variability

As well as changes in the morphology of the jet, variability in flux density can clearly be seen looking at the spectra in Figure 4.5 and Figure 4.6. It is immediately obvious that there is a high degree of variability in the jet emission of both sources, as the flux density at lower frequencies, where the jet emission dominates, varies significantly between the two epochs. On the other hand, the flux density at higher frequencies, where the disk emission dominates, is similar at both epochs, indicating that the disk emission does not vary significantly. In the N component in particular, the flux density has decreased significantly by a factor of ~ 5 compared with the data in Rodríguez et al. (1998). This suggests that the free-free emission from the jet has drastically decreased between the two epochs. The S component on the other hand, shows more modest although still significant variability between the two epochs with an increase in flux density by a factor of ~ 2 at 5 GHz. Some of the difference between the epochs may be due to differences in the synthesized beam of the observations, which were $0''.25 \times 0''.24$ and $0''.18 \times 0''.12$ in the Rodríguez et al. (1998) and Rodríguez et al. (2003) 8.3 GHz observations respectively, compared with $0''.20 \times 0''.18$ in the 5 GHz observation in this work. However, this is not enough to explain the decrease in the N component between 1998 and 2020. In the S component on the other hand, we see an increase in the flux density compared to 1998, but the smaller synthesized beam in our observations would lead to less flux density being recovered, not more, which shows that the increase in flux density we see is real.

For L1551 NE, the flux densities obtained can be compared with the data from (Reipurth et al. 2002), observed in 2000, as shown in Figure 4.7 and Figure 4.8. The synthesised beam size is similar for both epochs, $0''.37 \times 0''.30$ in our observations vs

$0''.34 \times 0''.27$ in Reipurth et al. (2002), and so they should recover similar scales of emission. While Reipurth et al. (2002) only observed L1551 NE at one frequency, it can still be seen from looking at Figure 4.7 and Figure 4.8 that the flux density at low frequencies seems to have remained approximately the same in component A while it has decreased in component B by a factor of ~ 2 between 2000 and 2021. This indicates that the free-free emission from the jet has decreased in component B.

This level of flux variability, particularly for the N component of L1551 IRS 5, where the free-free emission seems to have almost disappeared, is quite remarkable. For most YSOs observed, the free-free emission shows only modest variability on the order of 10 – 20% (e.g. Rodríguez et al. 2008, 2014; Carrasco-González et al. 2012). Although a few cases of more extreme variability have been seen previously. For example, the flux density of the radio jet in B335 was found to increase by a factor of $\gtrsim 5$ between 1994 and 2000 (Avila et al. 2001; Reipurth et al. 2002) while DG Tau A showed an increase in flux density by a factor of ~ 2 between 1994 and 1996 (Rodríguez et al. 2012).

Observations of jets and outflows on larger scales also show signs of variability in the ejection rate. Optical and infrared observations show that atomic and molecular jets contain closely-spaced knots near the star (< 0.1 pc) and larger, more distantly-spaced bows at further distances (Frank et al. 2014). These are thought to be caused by variations in the velocity of the jet material, which leads to internal shocks within the jet due to faster material collides with slower material ahead of it. It is not surprising then that we also see variability in the free-free emission from the ionized jet since it is likely tracing the same changes in mass-loss rate or velocity of the jet that is seen at larger scales.

Given the connection between accretion and ejection in YSO jets, it is expected that there should be a correlation between the thermal free-free emission, which traces the ionized jet, and accretion-tracing emission in the infrared and millimeter. Such a correlation has been seen for another outflow tracer, the flux density of the O I jet, which is well correlated with the mid-infrared excess of the inner disk and the optical excess of the hot accretion layer. However, it is not clear if there is such a correlation

between radio continuum emission and accretion tracers. For very young stellar objects, there appears to be a correlation between their radio continuum luminosity and their bolometric luminosity, which is mainly due to accretion (Anglada et al. 2018). On the other hand, the Class 0 object HOPS 383 showed a large increase in mid-infrared and bolometric luminosity, indicating an accretion burst, but showed no significant variability in flux density in the radio continuum at several epochs before and after the burst.

Given the strong variability seen in the radio continuum flux density seen in both L1551 IRS 5 and L1551 NE, they would be good candidates for future observations, along with other YSOs showing significant variability in free-free emission, to attempt to detect correlations between variability in free-free radio emission and in accretion-tracing emission.

4.3.4 Ionized Mass Loss Rate

From the fit to the spectra of the sources, the ionized mass loss rates of the jets can be estimated. From Reynolds (1986), the launching radius and ionized mass loss rate are given by:

$$\left(\frac{r_0}{\text{au}}\right) = 19.73 \left(\left(\frac{S_\nu}{\text{mJy}}\right) \left(\frac{\nu}{\text{GHz}}\right)^{-0.25}\right)^{1/2} \theta_0^{-1/2} \left(\frac{\nu_m}{\text{GHz}}\right)^{-7/8} \times \left(\frac{D}{\text{pc}}\right) \left(\frac{T}{\text{K}}\right)^{-1/2} (\sin i)^{-1/2} \quad (4.3)$$

$$\left(\frac{\dot{M}_{\text{ion}}}{\text{M}_\odot \text{ yr}^{-1}}\right) = 6.33 \times 10^{-19} \left(\left(\frac{S_\nu}{\text{mJy}}\right) \left(\frac{\nu}{\text{GHz}}\right)^{-0.25}\right)^{3/4} \theta^{3/4} \times \left(\frac{v}{\text{cm s}^{-1}}\right) \left(\frac{D}{\text{pc}}\right)^{3/2} \left(\frac{\nu_m}{\text{GHz}}\right)^{-0.26} \left(\frac{T}{\text{K}}\right)^{-0.075} (\sin(i))^{-1/4} \quad (4.4)$$

From this, it can then be shown that the ionized mass loss rate can be estimated from the spectrum using:

$$\begin{aligned} \left(\frac{\dot{M}_{\text{ion}}}{M_{\odot} \text{ yr}^{-1}} \right) &= 2.61 \times 10^{-19} K_{\text{ff}}^{0.6} \theta_0^{0.9} \left(\frac{D}{\text{pc}} \right) \left(\frac{T}{\text{K}} \right)^{0.075} \\ &\times \left(\frac{r_0}{\text{au}} \right)^{0.3} \left(\frac{v}{\text{cm s}^{-1}} \right) (\sin i)^{-0.1} \end{aligned} \quad (4.5)$$

For this work, an electron temperature of $T_e \sim 10^4$ K was assumed for all of the sources. It should be noted that while the mass loss rate derived depends on the value of T_e assumed, changing the value of T_e does not significantly affect the result as $\dot{M}_{\text{ion}}$ is only weakly dependent on T_e ($\dot{M}_{\text{ion}} \propto T_e^{0.075}$). A rough estimate for the opening angle can be found from the deconvolved major and minor axis of the jet emission: $\theta_0 \approx 2 \tan^{-1}(\theta_{\text{min}}/\theta_{\text{maj}})$. For L1551 IRS 5, using the values of θ_{min} and θ_{maj} obtained by Rodríguez et al. (2003), this gives estimates of $\theta_0 \approx 43^\circ$ and $\theta_0 \approx 28^\circ$ for the N and S component respectively. Unfortunately, neither our observations nor the observations from Reipurth et al. (2002) have sufficient resolution to estimate the major and minor axes of the jet. Therefore, we use a reasonable estimate of $\theta_0 \approx 30^\circ$ for these jets.

In addition to getting the launching radius and ionized mass loss rate based on our data, we can fit the data from Rodríguez et al. (1998) with the same model to get the values for a different epoch and compare it with the epoch of this work.

For the N component, the ionized mass loss rate estimated based on our data was:

$$\left(\frac{\dot{M}_{\text{ion}}}{M_{\odot} \text{ yr}^{-1}} \right) = 3.50 \times 10^{-10} \left(\frac{v}{200 \text{ km s}^{-1}} \right) \left(\frac{r_0}{\text{au}} \right)^{0.3}$$

while based on the Rodríguez et al. (1998):

$$\left(\frac{\dot{M}_{\text{ion}}}{M_{\odot} \text{ yr}^{-1}} \right) = 2.04 \times 10^{-9} \left(\frac{v}{200 \text{ km s}^{-1}} \right) \left(\frac{r_0}{\text{au}} \right)^{0.3}$$

This implies a factor of ~ 6 decrease in the ionized mass loss rate of the jet between 1998 and 2020.

For the S component, the ionized mass loss rate estimated based on our data was:

$$\left(\frac{\dot{M}_{\text{ion}}}{\text{M}_{\odot} \text{ yr}^{-1}} \right) = 1.84 \times 10^{-9} \left(\frac{v}{200 \text{ km s}^{-1}} \right) \left(\frac{r_0}{\text{au}} \right)^{0.3}$$

while based on the Rodríguez et al. (1998) data it was:

$$\left(\frac{\dot{M}_{\text{ion}}}{\text{M}_{\odot} \text{ yr}^{-1}} \right) = 1.30 \times 10^{-9} \left(\frac{v}{200 \text{ km s}^{-1}} \right) \left(\frac{r_0}{\text{au}} \right)^{0.3}$$

This suggests that there has been an increase in the ionized mass loss rate of the jet between the two epochs. Such changes are of course to be expected based on the change in flux density.

We can also compare the values of $\dot{M}_{\text{ion}}$ estimated with the total mass loss rate of L1551 IRS 5 to obtain a value for the ionization fraction of the jet. Based on O I observations with SOFIA from 2018 and 2020, Yang et al. (2022) calculated the mass loss rate of the outflow of L1551 IRS 5 to be $(1.3 \pm 0.8) \times 10^{-6} \text{ M}_{\odot} \text{ yr}^{-1}$. Unfortunately, the two individual jets are unresolved and so we only have the total mass loss rate for the whole system. But using this value as an upper limit for the total mass loss rate in each jet gives an estimate for the ionization fraction in each jet of $\lesssim 0.1\%$.

The same analysis can be performed for the jets of the two components of L1551 NE. For component A, the ionized mass-loss rate found based on the fit to the spectrum was:

$$\left(\frac{\dot{M}_{\text{ion}}}{\text{M}_{\odot} \text{ yr}^{-1}} \right) = 2.65 \times 10^{-10} \left(\frac{v}{200 \text{ km s}^{-1}} \right) \left(\frac{r_0}{\text{au}} \right)^{0.3} \left(\frac{\theta_0}{0.52} \right)^{0.9}$$

while for component B the ionized mass-loss rate found was:

$$\left(\frac{\dot{M}_{\text{ion}}}{\text{M}_{\odot} \text{ yr}^{-1}} \right) = 1.77 \times 10^{-10} \left(\frac{v}{200 \text{ km s}^{-1}} \right) \left(\frac{r_0}{\text{au}} \right)^{0.3} \left(\frac{\theta_0}{0.52} \right)^{0.9}$$

Unfortunately, we cannot calculate the previous ionized mass loss rate for L1551 NE from a fit to the spectrum as Reipurth et al. (2002) only observed the source at one frequency. However, from this measurement there appears to have been a decrease in flux between 2000 and 2021 in the free-free portion of the spectrum. Based on this,

we can infer that there has likely been a decrease in the free-free emission of the jet and therefore a decrease in the ionized mass loss rate. Similarly for component A, the free-free emission appears to be similar at both epochs and therefore the ionized mass loss rate has likely remained the same.

4.3.5 Radiative Transfer Modelling

In order to better understand the launching and collimation of the jet, radiative transfer modelling was performed for the jet in the S component of L1551 IRS 5 using the package *sf3dmodels* (Izquierdo et al. 2018) to model the jet and RADMC-3D (Dullemond et al. 2012) to then perform the radiative transfer calculations. *sf3dmodels* creates a three-dimensional grid of the temperature, ionization fraction, and density of the jet based on the model of (Reynolds 1986). To create this model, values for the mass-loss rate $\dot{M}_{\text{ion}}$, the velocity of the jet v_0 , electron temperature T_e , injection radius r_0 , half-width at the base of the jet w_0 , and inclination angle i of the jet are required. The model of an isothermal, collimated, ionized jet from Reynolds (1986) was used again, with $\epsilon = \frac{2}{3}$, $q_n = -2\epsilon = -\frac{4}{3}$, and a constant temperature and velocity in the jet.

Once the physical models have been generated, RADMC-3D is used to solve the equation of radiative transfer for free-free emission to generate a synthetic image of the region at a given wavelength. For L1551 IRS 5, we simulated an image at 5 GHz, since at this frequency almost all of the emission should be free-free emission as opposed to dust emission and so it should be possible to replicate the observed image using only free-free emission. The model image created was convolved with the same synthesized beam as the observed image in order to compare the two. We aimed to recreate the jet morphology, i. e. the double lobe structure, and the flux density seen in the observed image as best we can.

When creating the model, an estimate for the jet velocity of $v_0 = 200 \text{ km s}^{-1}$ and a temperature of $T_e = 10^4 \text{ K}$ was used. The half-width w_0 was taken to be given by r_0 and the opening angle of the jet $\theta_0 = 2(w_0/r_0)$. The morphology of the jet in the image was then mainly adjusted using the input value for r_0 , while the flux density

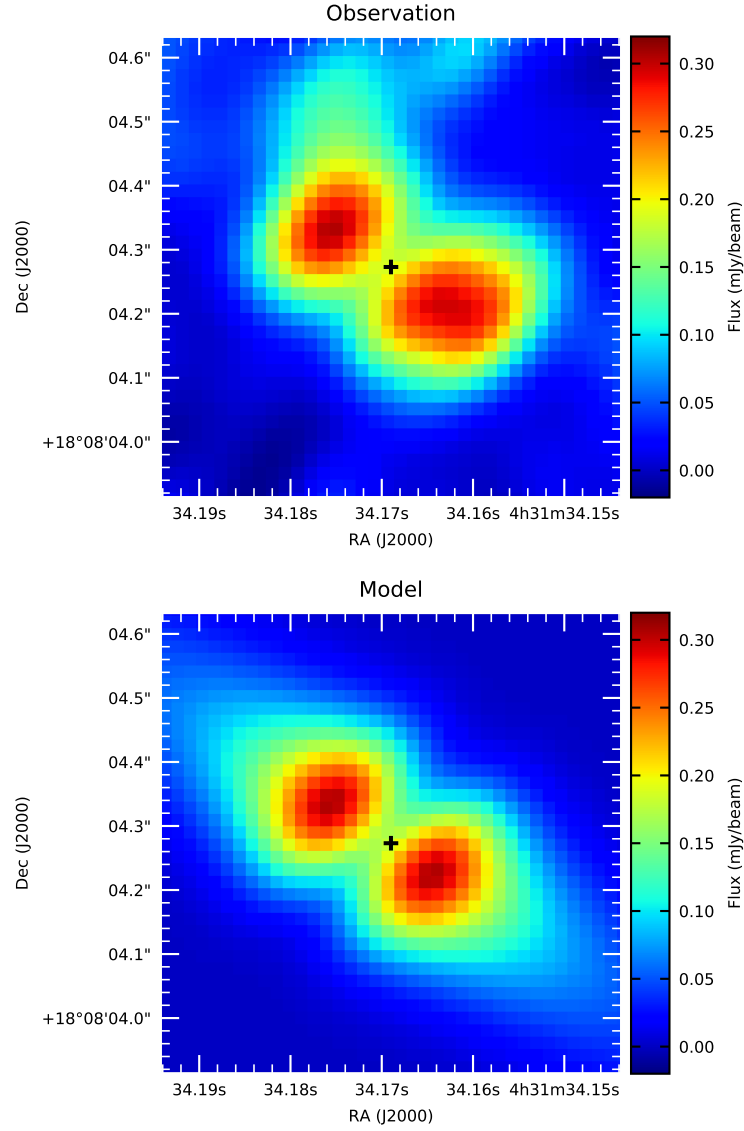


Figure 4.9: The VLA+e-MERLIN 5 GHz image of the S component of L1551 IRS 5 (top) compared with the image obtained from radiative transfer modelling (bottom). The peak of the emission in the 2 cm image is indicated by the black cross to show the position of the central star.

was adjusted using the value for $\dot{M}_{\text{ion}}$. It was found that the observed jet was best reproduced using values of $r_0 \approx 10$ au and $\dot{M}_{\text{ion}} \approx 4.7 \times 10^{-9} \text{ M}_{\odot} \text{ yr}^{-1}$. This can be seen and compared with the observed image in Figure 4.9. Note that the ionized mass loss rate obtained is similar to that predicted by Equation 4.5 for $r_0 = 10$ au, which would be $\dot{M}_{\text{ion}} = 3.7 \times 10^{-9} \text{ M}_{\odot} \text{ yr}^{-1}$. Assuming that the jet is collimated at scales $\sim r_0$, these results suggest collimation on scales of ~ 10 au, consistent with previous estimates for the collimation distance in low-mass YSO jets (Frank et al. 2014).

4.4 Conclusions

L1551 IRS 5 and L1551 NE were successfully imaged at a wide range of frequencies using high spatial resolution observations with e-MERLIN, the VLA, and ALMA. In particular, the 5 GHz image of L1551 IRS 5 that was obtained is one of the highest resolution images of a radio jet from a low-mass YSO obtained to date, with emission resolved on a scale of only ~ 15 au. Emission was detected from all of the components of the two binary systems at all frequencies. By plotting and fitting the spectra of the sources, it is apparent that free-free emission dominates at lower frequencies while dust emission dominates at higher frequencies, as expected. Although for the N component of L1551 IRS 5, there appears to be very little free-free emission present.

High levels of variability were detected from both components of L1551 IRS 5 and one of the components of L1551 NE, in both the flux density and the morphology of the jets. In particular, the N component of L1551 IRS 5 appears to decline in flux density by a factor of ~ 5 at lower frequencies, suggesting that the free-free emission of this source has almost disappeared.

From fitting the spectra, estimates for the ionized mass loss rates of each source were derived. By comparing the values obtained to an estimate for the total mass loss rate based on atomic line observations, it was estimated that the ionization fractions of the jets in L1551 IRS 5 are on the order of $\sim 0.1\%$.

Given the variability in flux density of the free-free emission of the jets, this implied there is variability in their ionized mass loss rates. By comparing the mass loss rates

derived based on the data in this work and from a previous epoch, it was shown that there is significant variation in the ionized mass loss rate of the jets, particularly the N component of L1551 IRS 5, on timescales of several years.

Based on the significant variability seen in the free-free emission in the two sources in this work as well as in other YSOs, it is important that multi-frequency observations of YSOs at cm-wavelengths should ideally be as close in time as possible. Otherwise there may be issues when comparing flux densities at different frequencies if they are at significantly different epochs.

The variability seen in L1551 IRS 5 and L1551 NE would make them good candidates for investigating a possible correlation between the free-free radio emission and accretion-tracing emission in YSOs, such as their mid-infrared or bolometric luminosities. These objects, along with other YSOs showing strong variability in their free-free emission, could be monitored over several years at different epochs to see if any such correlation can be detected.

The high resolution and sensitivity achieved in the 5 GHz image of L1551 IRS 5 made it possible to study the launching and collimation zone of the jet of the S component in detail. By creating a physical model and then using radiative transfer modelling to create simulated image which could be compared with the observed image, it was possible to obtain estimates for the collimation distance and mass loss rate of the jet. This suggested that the jet is collimated on scales of $\lesssim 10$ au, similar to previous upper limits for the collimation distances of YSO jets.

These results show the use of high-resolution, high-sensitivity observations for the study of radio jets. Future ultra-sensitive interferometers such as the Square Kilometer Array (SKA) or the Next-Generation Very Large Array (ngVLA) will allow for many more YSOs to be studied in this way and potentially help us learn more about the innermost regions of protostellar jets.

Chapter 5

Conclusions and Future Work

The main goal of this thesis was to explore ways of learning more about YSOs and star formation using radio astronomy. I showed that low-frequency observations can unlock new ways of studying YSOs, through the measurement of magnetic field strength in the jet using the Razin effect, and through the detection of coherent radio emission, previously rarely detected from YSOs. As well as this, at GHz-frequencies, I showed that high-sensitivity high-resolution observations allow us to study the inner region of the jet and the recently ejected material in detail. In the following section, I summarize and present the final conclusions on the main findings of this paper. I then comment on suggestions for future work in this field and what future instruments will allow us to do.

5.1 Summary and Conclusions

In chapter 2, I successfully detected synchrotron emission in two emission knots in the jet of the YSO DG Tau A using LOFAR, the second detection of a YSO and the first detection of non-thermal emission from a YSO at these frequencies. In one of the emission knots, a low-frequency turnover is observed in its spectrum, the first time such a turnover has been observed in non-thermal emission in a YSO jet. After considering the possible explanations for this turnover and fitting models for each of them to the spectrum, I determined that the Razin effect is the most likely explanation. From the

Razin effect fit, I was then able to measure the magnetic field in the jet, obtaining a value of $\sim 20 \mu\text{G}$ at a distance of $\sim 3 \times 10^3$ au from the star, a value that compares well with values obtained previously in other YSOs using different techniques. This potentially provides a new method of measuring magnetic field strengths in YSO jets, a quantity which has previously proven difficult to measure.

In chapter 3, I then showed further uses for low-frequency observations of YSOs, through the discovery of coherent radio emission bursts from the two YSOs KPNO-Tau 14 and LkCa 4. From the high brightness temperatures and polarization fractions of the bursts, it was determined that the emission must be due to electron-cyclotron maser (ECM) emission, a form of emission previously rarely seen from YSOs. This allowed the magnetic field strength to be measured near the surface of the star in both objects, given that the emission must have occurred at the local cyclotron frequency of the source. This gave a value of $40 - 70$ G. The ECM emission was determined to most likely be due to a breakdown of co-rotation between the plasma and the stellar magnetosphere. However, another exciting potential explanation is that it could be generated by a star-planet interaction, either through an interaction of the stellar wind with the planetary magnetosphere of an orbiting exoplanet, or through a sub-Alfvénic interaction between the stellar magnetosphere and an exoplanet, similar to the Jupiter-Io interaction in our own solar system. This could be an excellent new way of detecting exoplanets around YSOs, which have proven difficult to find with traditional detection methods due to the high stellar activity in YSOs.

In chapter 4, I then used high-resolution observations at a wide-range of frequencies to study the jets of two binary YSOs, L1551 IRS 5 and L1551 NE, in detail. This showed an unusually high level of variability in the free-free emission from both objects. In particular, one of the components of L1551 IRS 5 showed a factor of ~ 5 decrease in its flux density. By fitting the free-free emission and dust emission in the spectra of the sources, estimates were obtained for the ionized mass loss rates of the jets and by comparing this with previous observations, it was shown that there is significant variability in the ionized mass loss rates of the jets on the timescale of a few years. The

ionization fractions of the jets in L1551 IRS 5 were then estimated by comparison with the total mass loss rate obtained from atomic line observations and found to be on the order of $\sim 0.1\%$. Thanks to radiative transfer modelling of one of the components of L1551 IRS 5, the collimation distance of its jet was estimated to be $\lesssim 10$ au, in line with previous estimates. These results show the importance of high-resolution, high sensitivity observations of YSO jets. In particular, e-MERLIN allowed us to obtain some of the highest resolution images of the free-free emission of a YSO jet to date. Thanks to its high resolution, it will likely be a key instrument in this field for the foreseeable future.

5.1.1 Scales of emission

From the work in this thesis, it can be seen that studies of YSO radio emission allow us to study these objects on a range of scales and distances from the star. The coherent radio emission from KPNO-Tau 14 and LkCa 4 originates only a few stellar radii from the surface of the star. At the other extreme, the synchrotron emission in DG Tau A originates in a shock of the jet at ~ 3000 au from the central star, while the free-free emission from L1551 IRS 5 and L1551 NE traces emission from the base of the ionized jet at tens of au from the star.

For studies of YSO jets, combining observations of thermal free-free emission and non-thermal synchrotron emission allows us to study the properties of these jets both close to the launching region near the star, and at large distances along the jet axis. Studies of free-free emission help us to understand the launching mechanism causing these jets and give us estimates for the ionized mass-loss rate and ionization fractions of the jets. On the other hand, the presence of synchrotron emission at large distances from the central star shows that there are strong shocks present along the jet which accelerate particles to relativistic speeds. These are thought to be caused by material in the jet encountering dense ambient material in the surrounding molecular cloud, internal shocks within the jet due to variations in velocity, or as proposed for DG Tau A in this work and in Purser et al. (2018), due to dense material drifting into the jet.

5.1.2 Observations at different frequencies

An important overall conclusion that can be drawn from this work is the importance of studying YSOs at a wide range of frequencies, from low-frequency studies with LOFAR to mid and high frequencies with VLA, e-MERLIN, and ALMA, since they all reveal different aspects of these objects such as synchrotron emission from shocks in the jet at low frequencies (< 1 GHz), free-free emission from the ionized jet at mid frequencies ($\sim 1 - 20$ GHz), and dust emission from the disk at high frequencies (> 20 GHz).

The work in both chapter 2 and chapter 3 shows the use of low-frequency radio observations in studying YSOs as they have allowed us to observe previously unobserved or rarely observed phenomena. Little work had been done at these frequencies until previously, but the results both in this thesis and in other works such as the detection of thermal emission in T Tau (Coughlan et al. 2017), and the detection of synchrotron emission in the 300 - 700 MHz range from DG Tau A (Ainsworth et al. 2014) and HH80-81 (Vig et al. 2018), have given very interesting results indicating that more work should be done in this area.

In chapter 2 and chapter 4, we see how combining radio observations at different wavelengths can give us much more information than we can get from individual frequencies. For DG Tau A, combining the LOFAR data at 150 MHz with the GMRT data and VLA data at higher frequencies allowed us to see the turnover in the synchrotron emission and thereby estimate the magnetic field strength in the emission region after determining the emission was due to the Razin effect. In L1551 IRS 5 and L1551 NE on the other hand, the combination of e-MERLIN, VLA, and ALMA observations over a large range of frequencies of 5 - 350 GHz, allowed us to determine the contributions of free-free emission from the jet and dust emission from the disk to the flux density of the objects. This then allowed us to estimate the ionized mass loss rate for each of the jets and gave us the surprising result that there is almost no free-free emission present in the N component of L1551 IRS 5.

5.1.3 Magnetic Field

Magnetic fields are known to play an important role in the process of star formation, as they control accretion on to the star and the generation of outflows and therefore determine the final mass and angular momentum in the star and disk (Bouvier et al. 2007; Frank et al. 2014). In chapter 2 and chapter 3, we were able to study the magnetic fields in YSOs using methods not previously used in the study of YSOs.

While it is well-known that the launching of YSO jets involves the magnetic field of the star and/or disk, the strength and direction of the magnetic fields remain poorly understood, even though most other properties of YSO jets are well known (Ray 2009). Previously, the magnetic field direction has been measured from linear polarization measurements while the strength has been estimated using the equipartition method Carrasco-González et al. (2010). However, there are large uncertainties with the equipartition method since it is not certain if the source would be in equipartition. Therefore, the detection of the Razin effect in the synchrotron emission in knot C of DG Tau A provides a potentially very useful tool for measuring the magnetic field strength along the jet.

The structure of the magnetosphere in YSOs remains poorly understood, although in recent years much has been learned thanks to spectropolarimetric imaging at optical wavelengths (e.g. Donati et al. 2014; Hill et al. 2017). The detection of coherent radio bursts in KPNO-Tau 14 and LkCa 4 reveals important information about this topic, as it is a strong indication of magnetic activity in these objects. As well as this, the possibility that the emission is due to a star-planet interaction potentially could allow us to learn more about the planetary magnetosphere of orbiting exoplanets around YSOs.

5.2 Future Work

5.2.1 Low-frequency observations

Various kinds of emission have now been detected from YSOs using LOFAR, including free-free emission by Coughlan et al. (2017) and synchrotron emission and coherent radio emission in this work. An obvious goal for future work is then to try and detect these emission types at low frequencies in other sources.

Detections of synchrotron emission would be particularly useful since the detection of turnover in the spectrum due to the Razin effect would allow for magnetic field strength measurements within the jet. Note that a turnover may not necessarily occur at LOFAR frequencies. For example, if the source has either a higher electron density or a lower magnetic field strength the turnover would occur at higher frequencies. Therefore, it may be possible to detect a Razin effect turnover using the GMRT for example.

Unfortunately, DG Tau is currently the only low-mass YSO known to be associated with synchrotron emission. However, it has previously been detected in high-mass YSOs. The two best studied cases are Serpens SMM1 and HH80-81 (Rodríguez-Kamenetzky et al. 2016, 2017). HH80-81 in particular would be an interesting target to try and detect a turnover as the magnetic field strength was previously estimated from the equipartition method and this value could be compared with that from the Razin effect. In addition, the direction of the magnetic field is known thanks to the detection of linear polarization in the emission (Carrasco-González et al. 2010). HH80-81 was observed at frequencies of 300 - 1300 MHz using the GMRT (Vig et al. 2018), although it did not appear to show any turnover and so lower frequency observations are required.

As mentioned previously, the Taurus and Perseus Molecular Clouds were observed as part of the LOFAR Two-Metre Sky Survey (LoTSS). An important short-term goal will be to continue and complete the analysis of these regions. Taurus and Perseus each contain at least 414 and 369 YSOs respectively (Esplin et al. 2014; Young et al.

2015) and so they offer good chances of detecting further low-frequency YSO emission, from either free-free, synchrotron, or coherent radio emission.

When considering other star-forming regions to study, an important limitation is the distance to the region. Using the coherent emission from KPNO-Tau 14 as an example, the distance and flux density of the emission imply a brightness of $S_\nu D^2 = 81.22 \text{ Jy pc}^2$. Assuming an 8 hour LOFAR observation with an rms noise of $\sim 100 \mu\text{Jy beam}^{-1}$, coherent emission with this level of brightness could be detected at distances up to $\sim 400 \text{ pc}$. Star-forming regions in this range include the Taurus Molecular Cloud, Perseus Molecular Cloud, Ophiuchus Molecular Cloud, Serpens Molecular Cloud, and the Orion Molecular Cloud.

Unfortunately, apart from the Taurus and Perseus Molecular Clouds, there are few nearby star-forming regions observable with LOFAR. Due to the array’s northern latitude, regions at a declination $< -5^\circ$ are not observable with LOFAR, and so the Ophiuchus and Orion regions cannot be observed. As well as this, observations with LOFAR near the Galactic Plane are difficult due to increased noise. Therefore, observations of Serpens may be difficult. As a result, low-frequency studies of other star-forming regions may have to wait for the construction of the SKA (see subsection 5.2.4).

In the short-term, the Murchison Widefield Array (MWA) (Wayth et al. 2018) in Australia may be an option for studying star-forming regions in the southern sky, such as Ophiuchus and Orion. Unfortunately, due to its short baselines, MWA only has a maximum resolution of $\sim 1.5'$ at 150 MHz. This would likely impede any study of non-thermal emission in YSO jets. However, the detection of coherent radio emission may still be possible. In addition, coherent emission may occur at higher frequencies as well depending on the magnetic field strength in the source, potentially allowing detections with the GMRT or VLA.

5.2.2 High-resolution observations of YSO jets

In order to learn more about the launching and collimation of jets, it will be important to carry out further high-resolution observations of the inner regions ($< 100 \text{ au}$) of jets

in other YSOs to compare with the results presented here. e-MERLIN will likely be key in studying the inner regions of radio jets, at least in the short-term. Until the arrival of SKA2 and the ngVLA, it will probably remain the most high-resolution radio interferometer available at cm-wavelengths.

As well as studying the collimation of jets, high-resolution observations will be useful for studying the variability in their radio emission, particularly for seeing changes in the morphology of the jet. From variability in the flux density, we will also be able to monitor the variability in the ionized mass loss rates of these jets. A good idea would be to regularly observe the radio emission at cm-wavelengths from L1551 IRS 5 and other YSOs known to show variability in their free-free emission. This would give information on the timescales of the variability in the mass loss rates of these jets.

5.2.3 Radiative Transfer Modelling

While the radiative transfer modelling done in this work gives us important estimates on the collimation distance and ionized mass loss rate in a YSO jet, it is limited by the relatively simple model used for the jet. The Reynolds (1986) model assumes a constant mass loss rate and velocity. However, it is well known now that there can be significant variations in the accretion and therefore ejection rate of YSOs (Audard et al. 2014). This is shown by the presence of knots in the jet, caused by shocks due to velocity variations as faster moving material catches up with slower moving material (Ray & Ferreira 2021b). Therefore to improve the modelling of the radio emission of the jet, this should be accounted for. A software package known as Radio Jets in Python (RaJePy) is currently under development which can model variable mass loss rates in the jet (Purser, in prep.). Using this to perform the radiative transfer modelling for L1551 IRS 5 and other YSOs observed may give a more accurate and detailed model of the inner region of the jet.

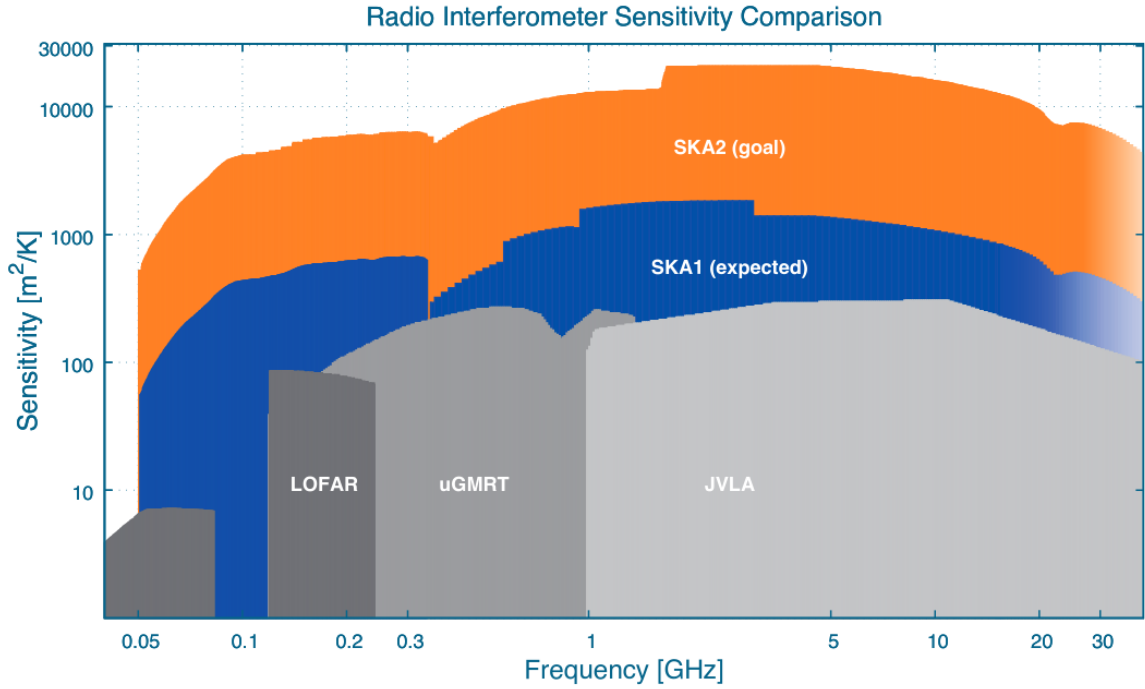


Figure 5.1: A comparison of the expected sensitivity of SKA1 and SKA2 with the current sensitivity of LOFAR, uGMRT, and the VLA. Credit: SKAO. Licensed under CC BY 3.0 (<https://creativecommons.org/licenses/by/3.0/>).

5.2.4 Future Instrumentation

In the next decade, radio astronomy is expected to undergo a revolution thanks to the construction of the Square Kilometer Array (SKA) and its unprecedented sensitivity. The first phase of construction (SKA1) is expected to be completed in the late 2020s and it will be comprised of two arrays: SKA1-Low located in Australia which covers 50 to 350 MHz, and SKA1-Mid, located in South Africa which covers 350 MHz to 15 GHz. It will be able to achieve angular resolutions of $5 - 10''$ with SKA1-Low and $0.05 - 0.5''$ with SKA1-Mid. In addition, it will achieve sensitivities up to an order of magnitude better than can be achieved with LOFAR and the VLA at the moment (see Figure 5.1).

SKA1-Low will significantly expand the possibilities for studying coherent emission from YSOs, allowing detections of fainter bursts and detections of bursts from star-forming regions further away, helping us to see how common this form of emission is from YSOs. With an order of magnitude increase in sensitivity compared to LOFAR, a coherent radio emission burst with the same luminosity as that of KPNO-Tau 14

could be detected at distances up to ~ 1300 pc. The improved sensitivity will also increase the chances of detecting synchrotron emission at low frequencies and therefore potentially allowing the detection of a Razin effect turnover in other YSOs.

With SKA1-Mid, it will be possible to carry out much more sensitive studies of radio jets in YSOs at high resolutions. This will allow us to study the base of the ionized jet in detail, helping us to identify the launching mechanism responsible (Anglada et al. 2015). The detection of radio recombination lines could also allow us to study the 3D kinematics of the jet at very small scales. For nonthermal emission, SKA1 also offers the possibility of further detections of linear polarization in YSO jets, similar to Carrasco-González et al. (2010), allowing us to study the structure of the magnetic field within YSO jets. Once the second phase of construction (SKA2) is completed, these capabilities will be further enhanced. This is expected to expand the array outside of South Africa with baselines 1000s of km in length, giving unprecedented resolution and sensitivity, allowing us to resolve in detail the launching and collimation zones of YSO jets.

As well as the SKA, the Next Generation Very Large Array (ngVLA) is expected to dramatically increase our capabilities for radio observations at GHz frequencies. With a planned maximum baseline of ~ 1000 km, it will be able to observe with a resolution of ~ 1 mas, making it a superb instrument for carrying out high-resolution studies of YSOs. For nearby star-forming regions (< 500 pc), this will allow studies of the emission on scales of only $\lesssim 0.5$ au. It will also be able to observe at higher frequencies than SKA, with a total frequency range of $1 - 116$ GHz. This will allow us to carry out high-resolution studies of the dust emission at these frequencies, which can then be compared with the free-free emission at lower frequencies.

Appendix A

Light curves of background field sources

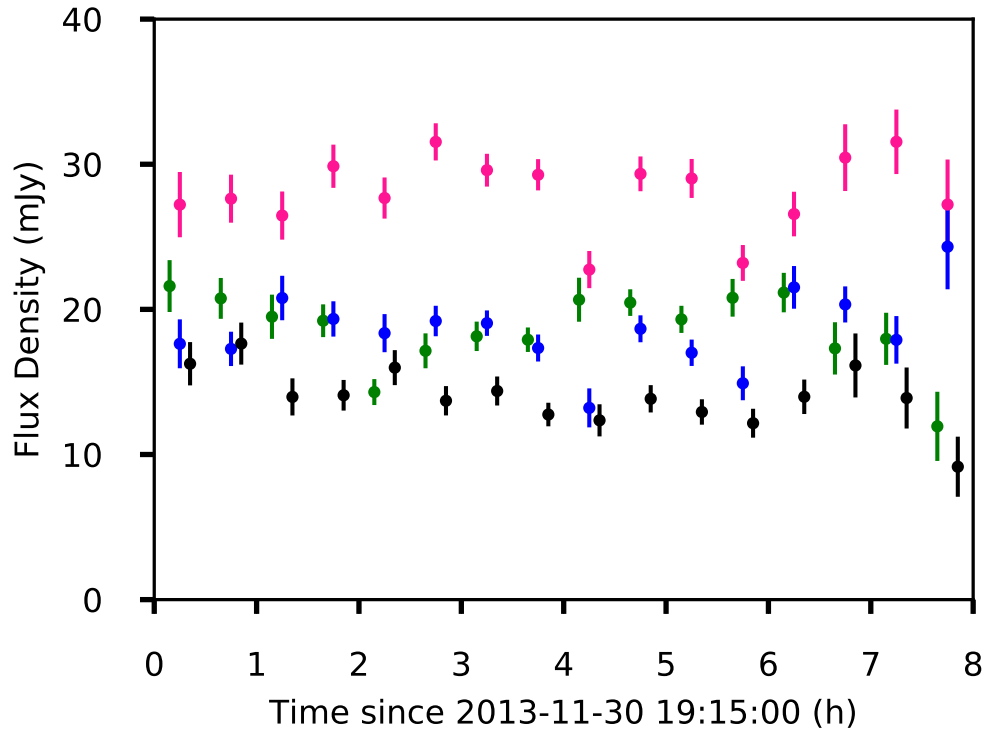


Figure A.1: Light curves of the sources in the field of KPNO-Tau 14: J043314+261818 (pink), J043316+261813 (blue), J043349+260753 (green), and J043331+260704 (black). Note that the fluxes of J043349+260753 and J043331+260704 are offset by -0.1 h and +0.1 h, respectively, for readability. The error bars are given by the fitting error σ_{fit} given by PyBDSF and the root-mean-square noise σ_{rms} added in quadrature: $\sigma = \sqrt{\sigma_{\text{fit}}^2 + \sigma_{\text{rms}}^2}$.

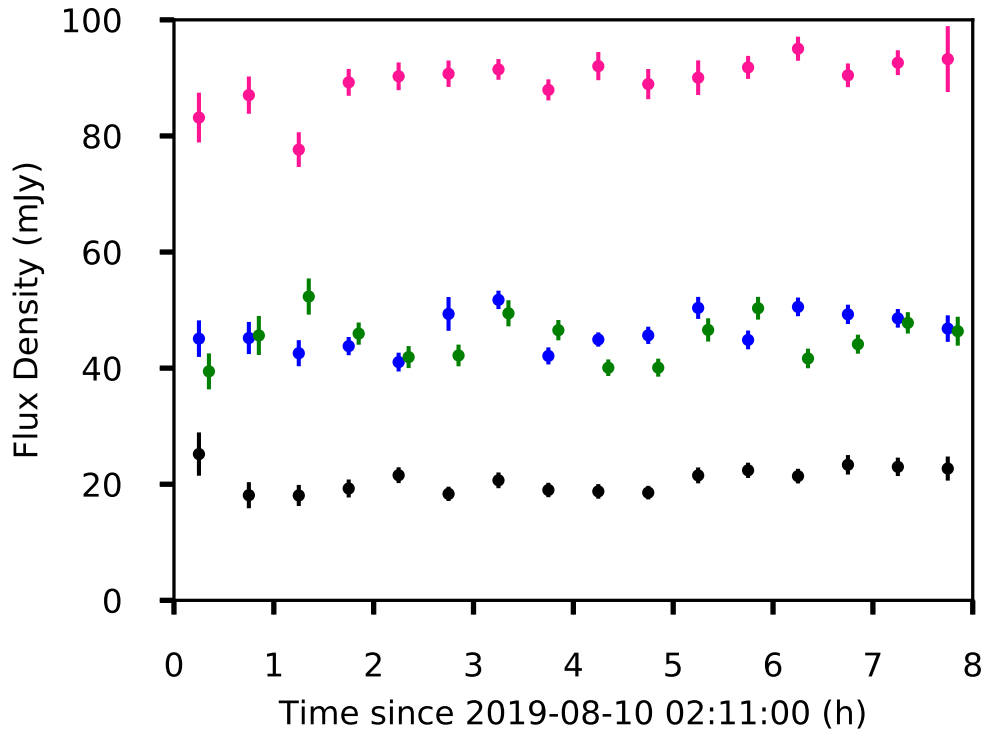


Figure A.2: Light curves of the sources in the field of LkCa 4: J041556+281645 (pink), J041732+281048 (blue), J041640+280521 (green), and J041607+281833 (black). Note that the fluxes of J041640+280521 are offset by +0.1 h for readability. The error bars are given by the fitting error σ_{fit} given by PyBDSF and the root-mean-square noise σ_{rms} added in quadrature: $\sigma = \sqrt{\sigma_{\text{fit}}^2 + \sigma_{\text{rms}}^2}$.

Appendix B

Flux density measurements of KPNO-Tau 14 and LkCa 4

Table B.1: Flux densities of KPNO-Tau 14 in Stokes I and V that were used in the light curve in Fig. 3.2.

Time (hrs)	Stokes I (mJy)	Stokes V (mJy)	Pol. Fraction (%)	Beam
0 - 0.5	≤ 4.26	≤ 3.60	-	$20''.3 \times 3''.2, -48^\circ$
0.5 - 1.0	≤ 3.21	≤ 2.85	-	$18''.7 \times 3''.1, -52^\circ$
1.0 - 1.5	≤ 3.44	≤ 2.88	-	$16''.6 \times 3''.1, -57^\circ$
1.5 - 2.0	11.25 ± 1.15	10.14 ± 0.82	88^{+8}_{-9}	$15''.0 \times 3''.2, -62^\circ$
2.0 - 2.5	≤ 2.70	≤ 2.13	-	$15''.0 \times 2''.9, -68^\circ$
2.5 - 3.0	4.16 ± 0.87	2.43 ± 0.69	57^{+20}_{-18}	$13''.8 \times 3''.1, -73^\circ$
3.0 - 3.5	4.24 ± 0.87	5.19 ± 0.67	91^{+7}_{-10}	$13''.0 \times 3''.4, -82^\circ$
3.5 - 4.0	8.70 ± 0.85	8.38 ± 0.65	91^{+6}_{-8}	$14''.2 \times 4''.1, 83^\circ$
4.0 - 4.5	≤ 3.15	2.94 ± 0.65	≥ 93	$20''.4 \times 4''.2, 66^\circ$
4.5 - 5.0	≤ 2.55	≤ 1.89	-	$13''.4 \times 4''.2, 62^\circ$
5.0 - 5.5	≤ 2.55	≤ 1.89	-	$7''.4 \times 4''.2, 61^\circ$
5.5 - 6.0	3.88 ± 0.83	≤ 1.92	≤ 49	$6''.2 \times 4''.2, 60^\circ$
6.0 - 6.5	≤ 2.58	2.91 ± 0.61	≥ 112	$5''.7 \times 4''.3, 60^\circ$
6.5 - 7.0	3.80 ± 0.98	2.81 ± 0.66	69^{+18}_{-20}	$5''.5 \times 4''.3, 61^\circ$
7.0 - 7.5	≤ 3.75	≤ 2.40	-	$5''.5 \times 4''.3, 62^\circ$
7.5 - 8.0	≤ 5.91	≤ 2.82	-	$5''.8 \times 4''.3, 62^\circ$

Notes. The time in the first column refers to the time in hours after the start of the observation (November 30, 2013, at 19:15:00). The polarisation fraction is shown for the times where both Stokes I and V are detected. When only Stokes I or V are detected, 3σ upper or lower limits, respectively, are given for the polarisation fraction. The dimensions and position angle of the synthesised beam in the image at each time are shown in the final column.

Table B.2: Flux densities of KPNO-Tau 14 in Stokes I and V that were used in the light curve of the first burst in Fig. 3.3.

Time (hrs)	Stokes I (mJy)	Stokes V (mJy)	Pol. Fraction (%)	Beam
1.50 - 1.55	≤ 8.40	≤ 8.07	-	$18''.3 \times 3''.0, -60^\circ$
1.55 - 1.60	≤ 8.76	≤ 9.21	-	$18''.5 \times 3''.0, -60^\circ$
1.60 - 1.65	13.76 ± 2.77	≤ 7.98	≤ 58	$18''.2 \times 3''.0, -61^\circ$
1.65 - 1.70	17.81 ± 2.90	18.30 ± 2.61	88^{+8}_{-13}	$18''.1 \times 3''.0, -61^\circ$
1.70 - 1.75	28.15 ± 3.02	27.10 ± 2.58	90^{+7}_{-8}	$18''.0 \times 3''.0, -62^\circ$
1.75 - 1.80	20.67 ± 2.81	24.30 ± 2.58	93^{+5}_{-8}	$17''.9 \times 3''.0, -62^\circ$
1.80 - 1.85	11.06 ± 2.78	17.10 ± 2.52	92^{+6}_{-10}	$17''.9 \times 3''.0, -63^\circ$
1.85 - 1.90	≤ 7.86	≤ 7.62	-	$17''.5 \times 3''.0, -63^\circ$
1.90 - 1.95	≤ 7.89	≤ 7.95	-	$16''.6 \times 3''.2, -63^\circ$
1.95 - 2.00	≤ 9.42	≤ 8.48	-	$16''.8 \times 3''.1, -64^\circ$

Notes. The time in the first column refers to the time in hours after the start of the observation (November 30, 2013, at 19:15:00). The polarisation fraction is shown for the times where both Stokes I and V are detected. When only Stokes I or V are detected, 3σ upper or lower limits, respectively, are given for the polarisation fraction. The dimensions and position angle of the synthesised beam in the image at each time are shown in the final column.

Table B.3: Flux densities of LkCa 4 in Stokes I and V that were used in the light curve in Fig. 3.5.

Time (hrs)	Stokes I (mJy)	Stokes V (mJy)	Pol. Fraction (%)	Beam
0 - 0.5	≤ 4.38	≤ 3.21	-	$26''.8 \times 3''.5, -46^\circ$
0.5 - 1.0	≤ 3.24	≤ 2.61	-	$25''.1 \times 3''.5, -50^\circ$
1.0 - 1.5	4.99 ± 0.99	3.30 ± 0.70	64^{+18}_{-16}	$23''.8 \times 3''.3, -55^\circ$
1.5 - 2.0	4.30 ± 0.76	3.76 ± 0.61	80^{+13}_{-15}	$22''.7 \times 3''.1, -59^\circ$
2.0 - 2.5	2.76 ± 0.68	3.00 ± 0.56	84^{+11}_{-16}	$21''.7 \times 2''.9, -65^\circ$
2.5 - 3.0	≤ 1.89	≤ 1.47	-	$20''.8 \times 2''.9, -71^\circ$
3.0 - 3.5	≤ 2.01	≤ 1.47	-	$19''.6 \times 3''.0, -77^\circ$
3.5 - 4.0	≤ 1.89	≤ 1.44	-	$15''.2 \times 4''.9, 66^\circ$
4.0 - 4.5	≤ 1.80	≤ 1.50	-	$21''.2 \times 3''.8, 84^\circ$
4.5 - 5.0	≤ 1.74	≤ 1.44	-	$13''.4 \times 4''.0, 67^\circ$
5.0 - 5.5	≤ 1.74	≤ 1.47	-	$6''.2 \times 2''.3, -52^\circ$
5.5 - 6.0	2.28 ± 0.62	≤ 1.53	≤ 67	$9''.8 \times 3''.9, 62^\circ$
6.0 - 6.5	≤ 1.83	≤ 1.53	-	$7''.5 \times 3''.9, 61^\circ$
6.5 - 7.0	3.95 ± 0.60	2.40 ± 0.53	60^{+17}_{-15}	$6''.5 \times 4''.0, 60^\circ$
7.0 - 7.5	4.46 ± 0.65	3.90 ± 0.56	82^{+11}_{-13}	$6''.1 \times 4''.0, 59^\circ$
7.5 - 8.0	≤ 2.25	≤ 1.86	-	$6''.0 \times 4''.0, 57^\circ$

Notes. The time in the first column refers to the time in hours after the start of the observation (August 10, 2019, at 02:11:00). The polarisation fraction is shown for the times where both Stokes I and V are detected. When only Stokes I is detected, 3σ upper limits are given for the polarisation fraction. The dimensions and position angle of the synthesised beam in the image at each time are shown in the final column.

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