KRITTIKA SUMMER PROJECTS 2023

Exploring the Radio Sky

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1. The Basics

1.1 Why Radio Astronomy

- The sky at radio wavelengths is very different from the sky at visible wavelengths. Objects bright at visible wavelengths, such as stars, are not what dominate the emission in the radio sky.
- Using radio waves, we can detect the thermal continuum and spectral-line emission from objects too cold to produce visible light, permitting studies of the cold interstellar medium of our Galaxy and others, as well as the cosmic microwave background, the relic radiation from the early universe.

Figure 1.1: NGC1004 as observed by the WISE telescope in micrometer wavelengths

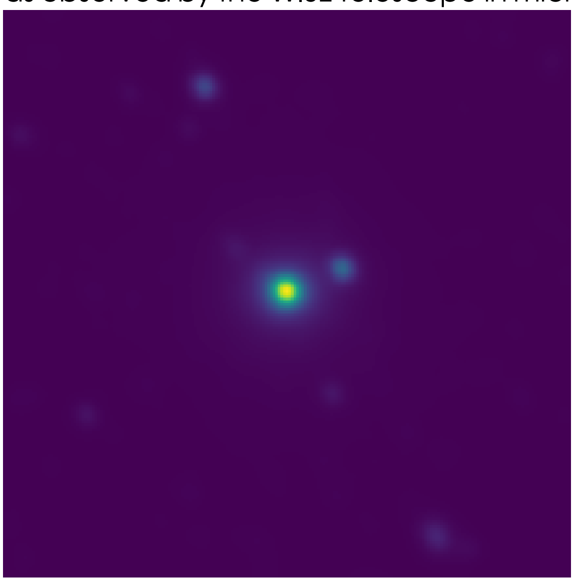
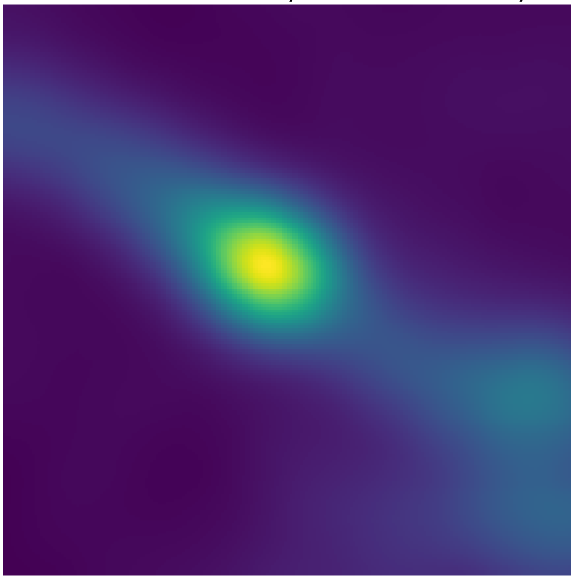


Figure 1.2: NGC1004 as observed by the NVSS survey in GHZ frequency



1.2 A Brief History of Radio Astronomy

For a long time we could not do radio astronomy due to the extremely weak nature of the astronomical radio waves reaching us. Radio astronomy was able to begin once we were able to aplify radio signals entering the antenna.
After this, radio astronomy made big leaps during the war due to the need of similar technology involved in radars
Over all this time, many astronomical radio sources were discovered like the galaxy centre, supernovae, sunspots, lightning and meteor showers.
The realization of a major avenue for study of large pieces of the universe using radio observations was initiated in 1944 when Jan Oort suggested to Hendrik van de Hulst that he calculate the wavelength of the emission line of hydrogen due to the spin flip of the electron. The calculations by van de Hulst predicted that this transition should emit radiation at 21 cm wavelength.
The first step in development of high resolution radio astronomy was taken when Sir Martin Ryle and D. D. Vonberg made the first astronomical observation using a pair of radio antennas as an interferometer.
Radio astronomy yielded discoveries of quasars, pulsars, and the cosmic microwave background, and continued advancing with the use of radio interferometers.
Nowadays, radio astronomy has reached the same level of operation as visible wavelength astronomy

1.3 Structure of a Traditional Radio Telescope

Radio Astronomy is very different from Visible Wavelength astronomy in a few ways:

- The Sun does not light up the whole sky at radio wavelengths
- At radio wavelengths, there is no scattering by the atmosphere
- At long radio wavelengths, observations can occur even with a cloud-covered sky

The Telescope design takes into account these unique situations

A Radio Telescope has 3 main parts

1.3.1 The Reflector

Most Radio Telescopes have a parabolic reflector, also known as the dish to collect and focus the radio light. As Radio waves are much easier to reflect than to refract, radio telescopes are not made using a refractor model
The radiation that reflects off the reflector can be collected at the focus, or be reflected again at the focus and collected behind the dish.
Some important aspects of the reflector are:

- The collecting area, which affects the sensitivity .
- The deviations in the surface of the reflector. The deviations must be much smaller than the wavelength of the light being collected
- The size of the reflector, which affects the resolution due to diffraction.

1.3.2 The Mount

This is the physical structure which holds and moves the dish with the help of motors. For the telescope to be able to point in any direction, the mount must have 2 degrees of freedom.
There are 2 types of mounts

1. Alt-Az mount
2. Equatorial mount

1.3.3 The Feed, Receiver and Computer

The dish focuses light from the sky to specially designed antennae called feeds. These antennae convert the EM radiation in free space to currents in a wire.
Each feed is connected to a receiver, which has 2 parts.

1. The front end, connected as close to the feed as possible, provides amplification and frequency conversion. The signal can then be sent a significant distance away using a coaxial cable.
2. The back end receives the signal from the cable. It contains a detector which measures the amount of power and converts it to a radio signal, which is then stored in a computer for future analysis.

A feed-receiver assembly is called a detector and is equivalent to a pixel in visible wavelength astronomy. Usually the number of detectors in a telescope is quite small. This is again in contrast to visible wavelength astronomy, where a single observation can give millions of pixels.

1.4 Exercise: Plotting the Jet Afterglow of the GW170817 event

GW170817 was a merger of two neutron stars, accompanied by gravitational and electromagnetic waves. For this activity, I considered the radiation to be having a non thermal emission, following the spectral index:

$$F_\nu = \nu^{-0.584}$$

at all frequencies. The lightcurve, which is the flux density (F_ν) as a function of time was calculated by me using two different sets of measurements.
This is explored in more detail in Section 5

Figure 1.3: 3GHz data points from the VLA telescope

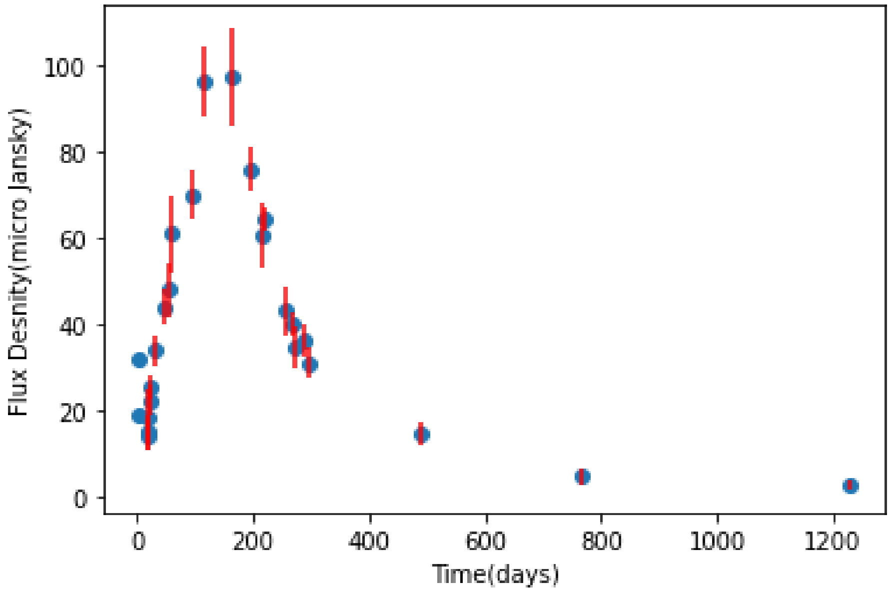
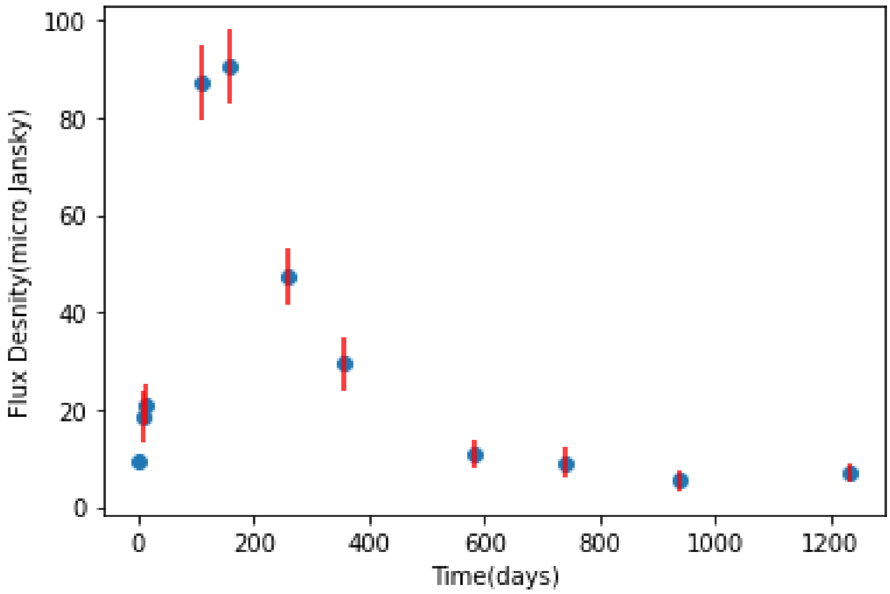


Figure 1.4: Data points from the Chandra telescope, normalized to 3GHz



2. Radiative Processes in Astronomy

2.1 Quantifying radiation

2.1.1 Total Energy

The total energy released by a source of a lifetime is an interesting yet very difficult to measure thing. It is not usually used while quantifying the radiation of a source.

2.1.2 Luminosity

Luminosity, also known as power, is the rate at which energy is emitted by the source. This is also not easy to measure as we do not receive all the power from the source. Only a very small portion of the energy emitted by the source reaches our telescope.

2.1.3 Flux

The energy collected by a telescope is influenced by the characteristics of the telescope, especially the collecting area of the telescope. All these effects are combined into a number called the effective area of the telescope. This flux can also be thought of in terms of normalizing for the distance from the source.
FLux helps us compare the observations from different telescopes that may have different areas.

$$F = \frac{L}{4\pi d^2} = \frac{P}{A_{eff}}$$

Where P is the power measured by the telescope

2.1.4 Flux density (wrt bandwidth)

Different telescopes are sensitive to different parts of the spectrum. Also, the emission from most sources is not uniform over the spectrum. Thus, Flux density, which is the flux at a given wavelength is a much more useful way to quantify the radiation of a source.

$$F_\nu = \frac{F}{\Delta\nu}$$
$$F_\lambda = \frac{F}{\Delta\lambda}$$

$$P = F_\nu A_{eff} \Delta\nu$$

2.1.5 Intensity

Intensity is the most fundamental way to quantify radiation as it is independent of the observer. Observers at all distances will agree upon the Intensity of a given source. This is because we normalize for the visible (angular) size of the source. Observers close to the source will see a brighter, but more spread out source and vice versa, and the ratio of the two will stay the same.

$$I_\nu = \frac{F_\nu}{\Omega}$$

where Ω is the angular size of the source. Intensity is, in fact, the correct description of the quantity that human eyes measure and the human brain interprets.

2.2 Black Body Radiation

Blackbody radiation is the emission of electromagnetic radiation from an object due to its temperature. All objects with nonzero temperature emit energy as photons, and the distribution of this radiation follows Planck's law. As the temperature increases, the emitted radiation shifts to shorter wavelengths.

2.2.1 The underlying process

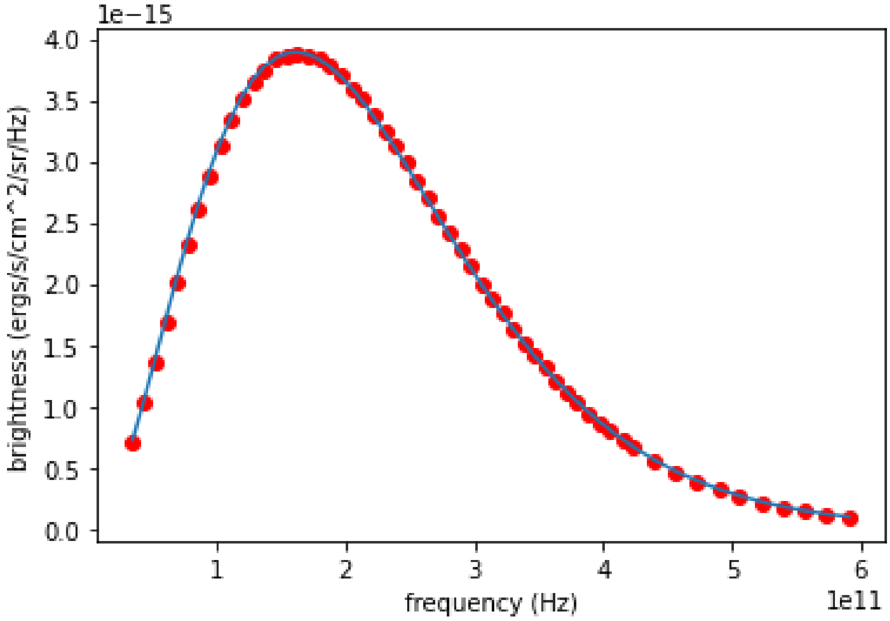
When a blackbody is in thermal equilibrium, the atoms or molecules inside it continuously absorb and emit photons. This fact with the quantization postulate gives rise to the Planck's blackbody radiation formula. This formula gives the spectral distribution of energy (intensity) as a function of wavelength or frequency and temperature.
Planck's law can be written in terms of frequency as well as wavelength, both of which give us the same information in slightly different ways.

$$B_\nu(\nu, T) = \frac{2h\nu^3}{c^2} \frac{1}{e^{\frac{h\nu}{k_B T}} - 1}$$
$$B_\lambda(\lambda, T) = \frac{2hc^3}{\lambda^5} \frac{1}{e^{\frac{hc}{\lambda k_B T}} - 1}$$

2.2.2 The Universe, an almost perfect black body

The Cosmic Microwave Background is a remnant of the big bang that allows us to probe the early stages of the universe. It is radiation that comes to us from every direction and is almost uniform.
Here, in Figure 2.1 I tried to fit a black body curve to this radiation and find the effective temperature.

Figure 2.1: Intensities of the CMB at different frequencies, fit to a black body curve at 2.74K



2.3 The 21cm line

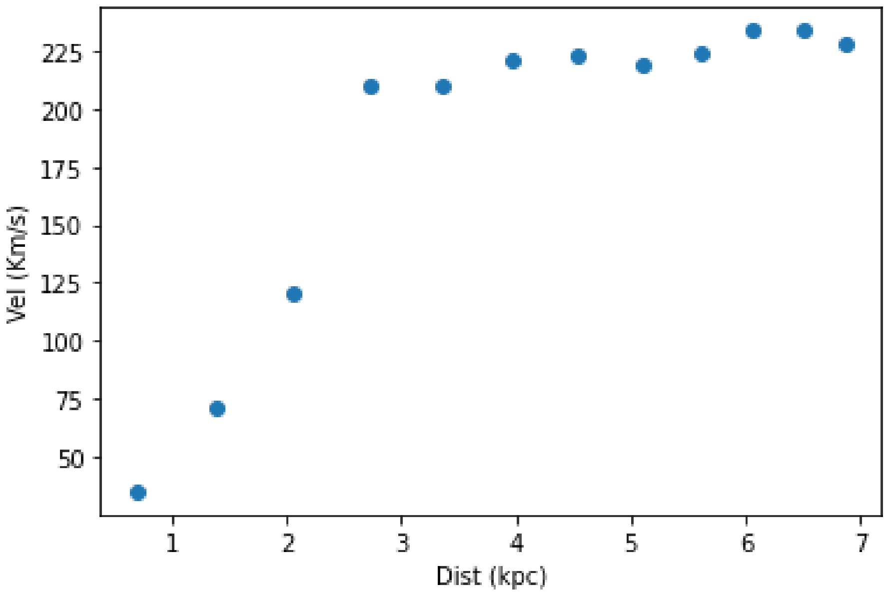
2.3.1 The underlying process

When an electron in a neutral hydrogen atom transitions from a higher energy state to the lowest energy state (the spin-flip transition), it releases energy in the form of a photon with a wavelength of 21 cm. This transition occurs due to the electron's spin.
This line is of great iterest to astronomers as it can penetrate the dusty regions of space that often obscure other wavelengths of light. It allows them to study the distribution and motion of neutral hydrogen gas in galaxies, as well as the large-scale structure of the universe.

2.3.2 Using the 21cm line to find the rotation curve of a galaxy

The 21cm line will be red-shifted when viewed from a frame in which the source has a radial component to its velocity.
Using this red-shift value, and the extremely thin nature of the 21cm line, we can calculate the speed of the source to a high accuracy.
I did the same in Figure 2.2

Figure 2.2: Rotation curve of a galaxy calculated using the 21cm line



3. Radio Telescopes

Unlike telescopes with deal with high frequency light, Radio Telescopes can not exploit the particle nature of light for their detectors, as the photons have insufficient energy to produce any effect on a semiconductor device. Instead, they detect radiation using large ensembles of photons and make use of the wave nature of light.

3.1 Reflectors

The dishes of most radio telescopes are parabolic reflectors. The parabolic shape causes all waves approaching the dish from the direction perpendicular to the entrance plane to come to a single point, known as the focus of the telescope.
If the waves are not exactly perpendicular, they still converge at a point, slightly offset from the focus (on the focal plane).
Having a receiver at this focus is usually inconvenient, so most telescopes have a secondary reflector which redirects the radiation to a point behind the primary reflector.
The primary reflector serves two important functions.

1. It collects and focuses the radiation from astronomical sources, making faint sources more detectable.
   The amount of radiation collected depends on the telescope's effective area (Aeff), which is closely related to the physical area of the primary reflector
2. It provides directivity, which is a telescope's ability to differentiate the emission from objects at different angular positions on the sky.

3.2 Beam Pattern

The beam pattern is a measure of the sensitivity of the telescope to incoming radio signals as a function of angle on the sky.

Because the sensitivity pattern is the same, whether the antenna receives or transmits—a principle known as the **reciprocity theorem** we are free to describe the pattern either way.
Ideally we would want the beam pattern to be extremely sharp, pointing only in 1 direction. But this is not possible due to diffraction.
Light coming from different parts of the sky interfere when they reach the receiver, and mix with each other. This pattern of constructive and destructive interference give rise to the beam pattern.
(i didnt include much of the maths)
For a telescope with a uniformly illuminated circular aperture, the total collected power is zero when the source is $1.22(\lambda/D)$ radians from the central axis.
As the off-axis angle increases further, the response goes through of peaks and valleys caused by partial constructive and destructive interference. These off-axis responses are called sidelobes and are undesirable as they can add confusion to observations.
The width of the central peak of the pattern is used to define the angular resolution of a single-dish telescope. By custom, we define this angular resolution as the full width at half maximum (FWHM) of the main beam of the telescope. This occurs (for a uniformly illuminated circular aperture) at $1.02(\lambda/D)$
Usually a small angular resolution is desirable, as it means that astronomical sources close together in angle on the sky can be distinguished

3.3 Feeds

Feeds are present at the centre of the dish and convert the EM waves into voltages. The feeds and receiver work well only for specific frequency ranges dictated by their design. Once again, Diffraction, determines the amount of power the feed collects and passes onto the receiver.
Ideally, we would like a feed-horn beam that had as close to uniform sensitivity out to the edge of the dish as possible, as this would yield a maximum response to the radiation from the source. However, we do not want the feed horn to detect contaminating back- ground radiation coming from beyond the edge of the dish
A quantity that describes how the feed horn's beam is distributed on the primary reflector is called the edge taper, which is defined as the ratio of the sensitivity at the center of the reflector to that at the edge.

3.4 Surface Errors

There are always manufacturing imperfections that limit its surface accuracy. We can characterize an imperfect reflector by the root mean square (rms) deviations, δ z
Such deviations will cause the path length to the focus to be slightly different for various parts of the reflector, and hence incomplete interference
The effect of surface errors on the collecting area is described by the Ruze equation, which is given by

$$A_\delta = A_0 e^{-(4\pi\delta z/\lambda)^2}$$

The rms error should be much smaller than the wevalength for the telescope to have a reasonable preformance

3.5 Noise and Temperature

All components in a radio telecope generate their own (unwanted) electrical signals. This interferes with our ability to observe radiation from an astronomical source.

3.5.1 Charachterising Noise

Nyquist in 1928, found that a resistor in the circuit will add electrical noise with a power per Hz that depends solely on the resistor's temperature. For this reason, the electronic power in a circuit, in general, can be described in terms of an equivalent temperature, T_{equiv} , which is equal to the temperature of a resistor that would produce the same amount of power as the resistor.
Following this convention, power is described using an equivalent temperature given by

$$T_{equiv} = \frac{P}{k\Delta\nu}$$

The equivalent temperature from the astronomical source is called the antenna temperature. The temperature due to noise from system components is called the noise temperature.

3.5.2 Gain

At each stage in a reciever, the signal is either amplified or reduced. We quantify this as the gain of that component. Each of these components also add some noise to the system.
An amplifier's noise temperature is defined by the equivalent temperature of the noise power as if it was introduced at the input to the amplifier, and hence it is amplified along with the astronomical signal.
So, if we have multiple amplifiers in succession, the net noise temperature is given by

$$T_N = T_{N1} + \frac{T_{N2}}{G_1} + \frac{T_{N3}}{G_1 G_2} \ldots.$$

This makes it very easy to compare the effects of different components on the noise.
Note that that first device the radiation enters into immediately after the feed, therefore, is the most critical in determining the total noise temperature. So this first amplifier should be a stateof-the-art device, and not a device that is mass-produced for commercial use.

3.5.3 Switched Measurements

For most sources, $T_A << T_N$, so even if we look at the blank sky, the power received is significant. Because of this, we cant make direct measurements of the sky.
We must make switched power measurements in which we measure the difference in voltage between when the telescope is aimed at the astronomical source (called an on-source observation) and when it is aimed at blank sky (called an off-source observation)
This removes the offset in measured power caused by noise, but it does not remove the fluctuations in noise. It is the fluctuations of the noise power that limits the ability to detect a weak astronomical source.

The variance in this noise(for low frequencies radiation like radio waves) is proportional to the number of photons per mode.

$$\sigma_P = \frac{P_N}{\sqrt{\Delta t \Delta \nu}}$$

3.6 Radio Frequency Interference

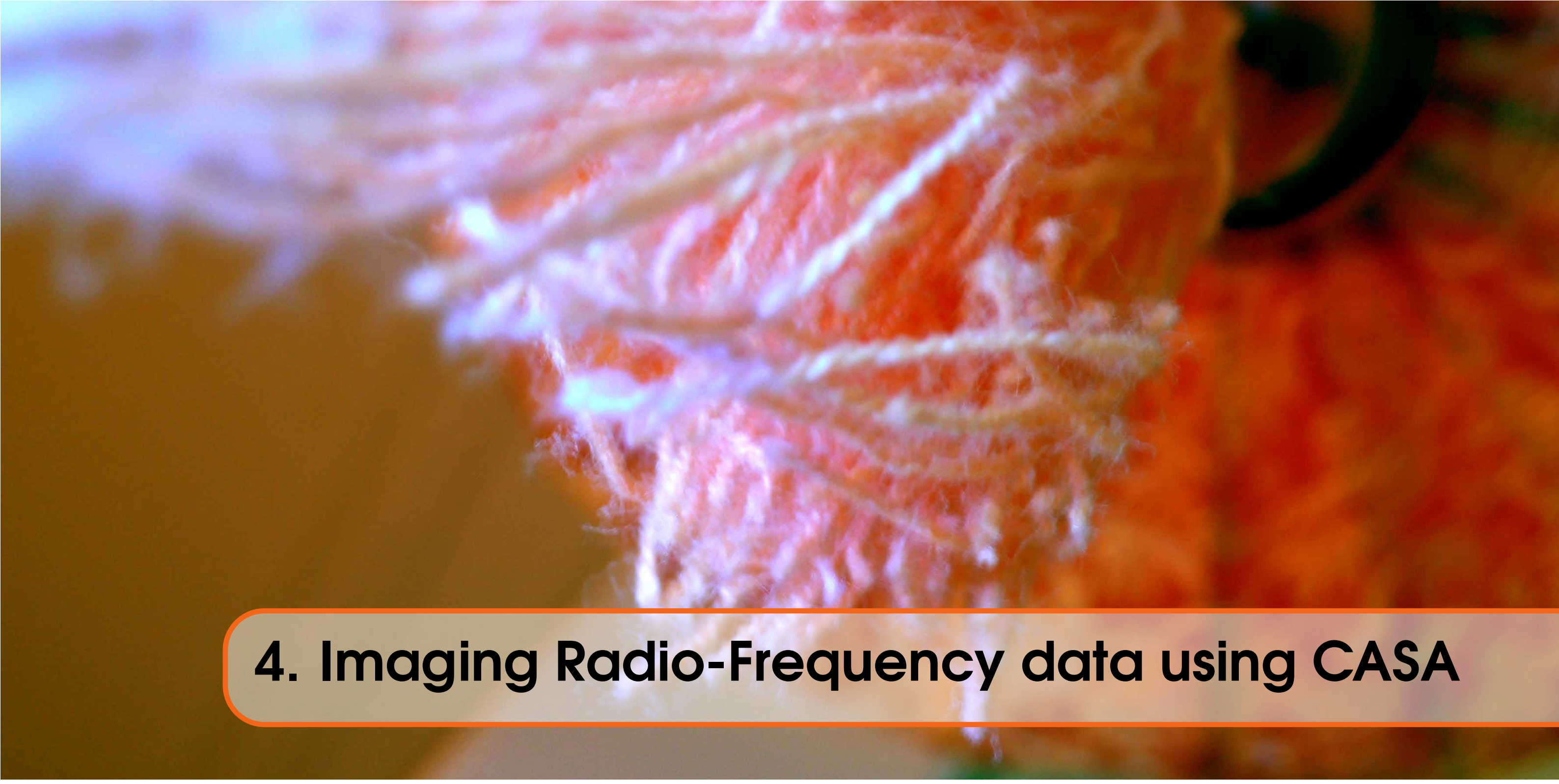
3.6.1 Sources

RFI is any radio signal, made by nature or by humans, that interferes with the radio waves we wish to detect. There are different sources of RFI at different frequencies. for example:

- GPS, WiFi at 0.8 - 2.5GHz
- satellites at 10-25GHz
- lightning at lower frequencies

3.6.2 Avoiding RFI

- locate the telescope as far away from human activity as is reasonably possible and to install filters to keep interfering signals out of the receiver path.
- divide the observed passband into many spectral channels, and remove the channels which have RFI



4. Imaging Radio-Frequency data using CASA

4.1 CASA

CASA, the Common Astronomy Software Applications package, is the primary data processing software for the Atacama Large Millimeter/submillimeter Array (ALMA) and NSF's Karl G. Jansky Very Large Array (VLA), and is frequently used also for other radio telescopes. The CASA software can process data from both single-dish and aperture-synthesis telescopes, and one of its core functionalities is to support the data reduction and imaging pipelines for ALMA, VLA and the VLA Sky Survey (VLASS).
I used CASA to image to start TW Hydra and the supernova 3C391 using the calibrated data.

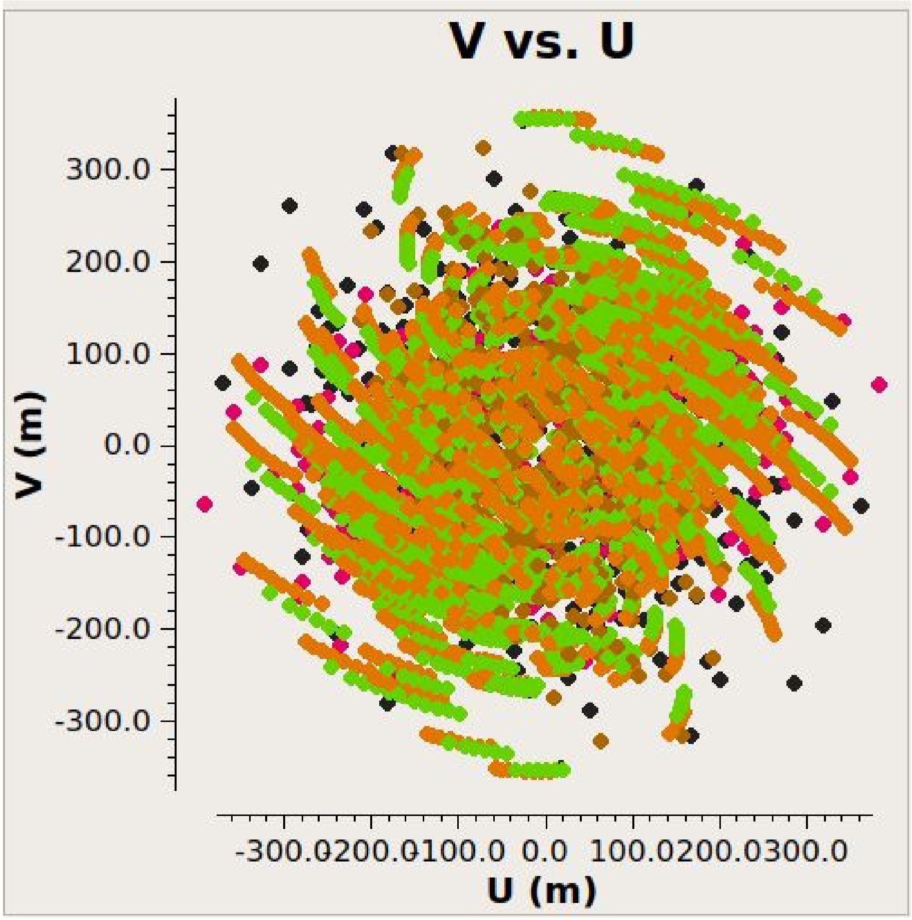
4.2 Interferometry

Interferometry is a technique which uses the interference of superimposed waves to extract information. A radio interferometer is an apparatus consisting of two or more separate antennas that receive radio waves from the same astronomical object and are joined to the same receiver. The antennas may be placed close together or thousands of kilometres apart. The displacement between a pair of antennae is called a baseline. Longer baselines allow us to capture more detailed images.

4.3 UV Coordinates

The UV coordinates correspond the intensities captured by different 2-dimensional baselines in a telescope array. Each baseline corresponds to a different spatial Fourier component on the final image.
The calibrated data is in the form of UV coordinates in a .ms file and can be viewed using the plotms function in CASA (Figure 4.1).

Figure 4.1: TW Hydra in UV coordinates



4.4 Sky Coordinates

Using the inverse fourier transform, these UV coordinates can be converted into sky coordinates. This gives rise to the dirty image. The dirty image has a lot of unwanted artefacts (Figure 4.2) as we have only a finitie number of fourier components using a finite number of antennae.

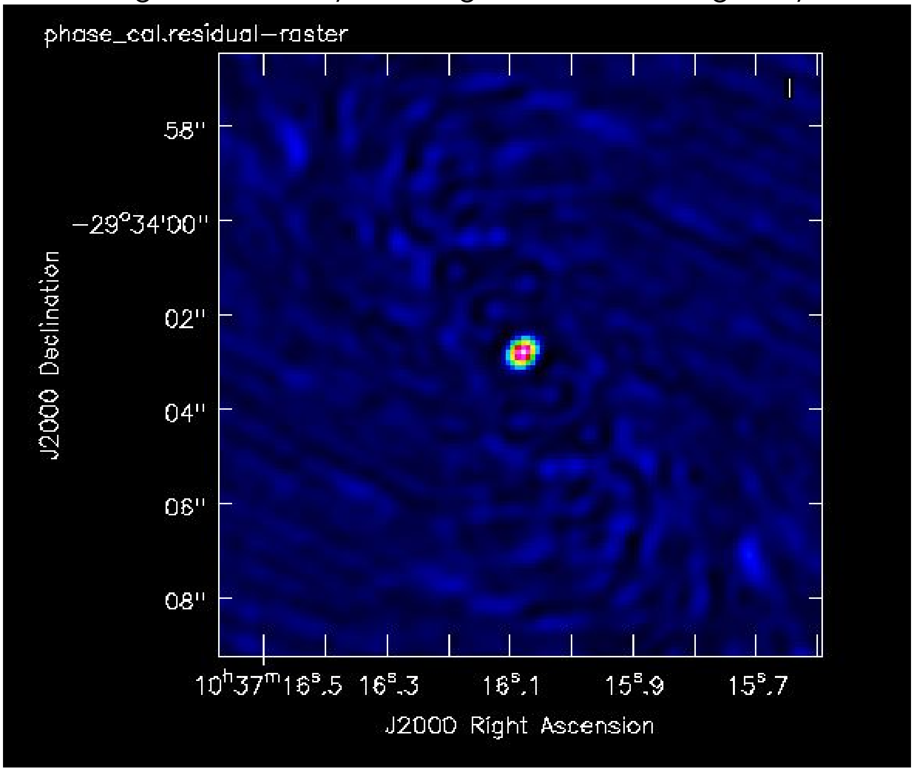
4.5 Cleaning the Image

Using CASA's tclean command we can clean this image and improve the SNR of the image. Here, we apply a mask on the image, and the clean algorithm performs a de-convolution on the masked part of the image, taking into account the fourier components we could measure.

As we can see, most of the artefacts from Figure 4.2 have been removed in Figure 4.4

Cleaning is an iterative process. One might have to do it a few thousand times to get good results. Cleaning should be done till one can no longer identify the source in the residual image (i.e. the residual reaches the same level as the noise)

Figure 4.2: TW Hydra Image before cleaning (dirty



4.6 Supernova 3C391

I did the same process on a supernova

Figure 4.3: TW Hydra Mask used for cleaning. The white area on this image was marked as the source and the deconvolution tries to recover this area.

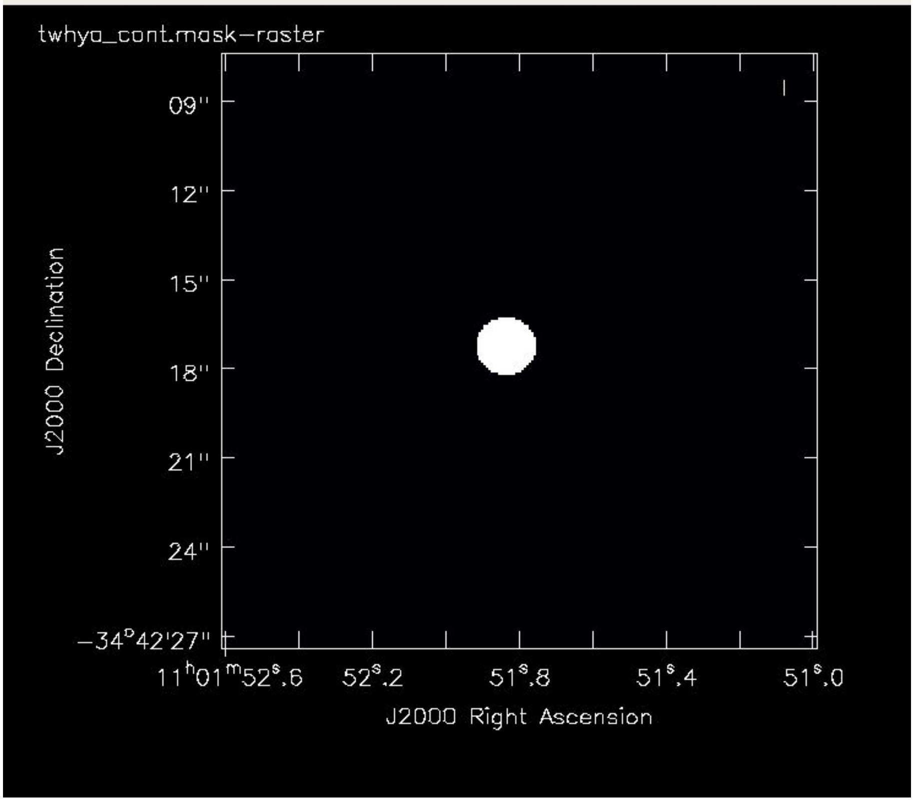


Figure 4.4: TW Hydra Image after deconvolution (clean)

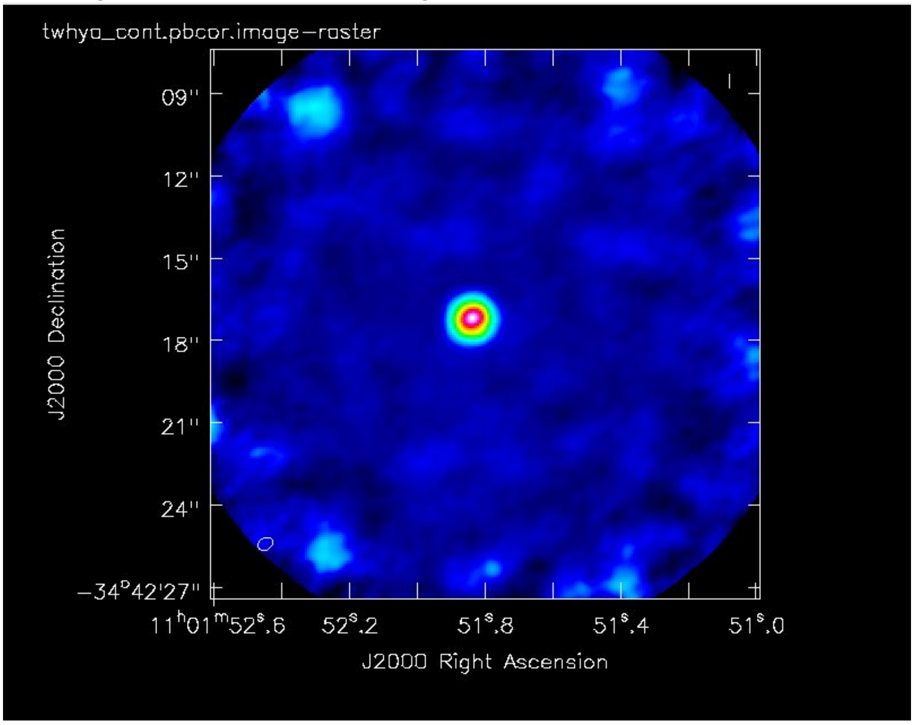


Figure 4.5: TW Hydra Residual. This is the radiation left after deconvolution. Notice that here, the source does not stand out from the surroundings.

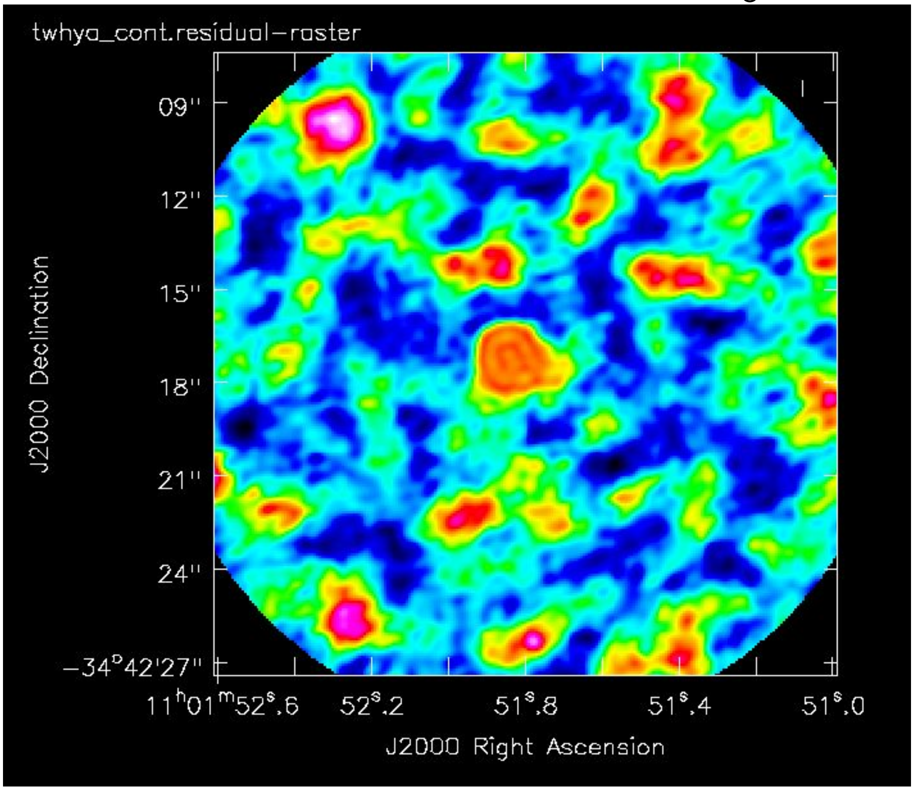


Figure 4.6: Supernova 3C391 Cleaned Image

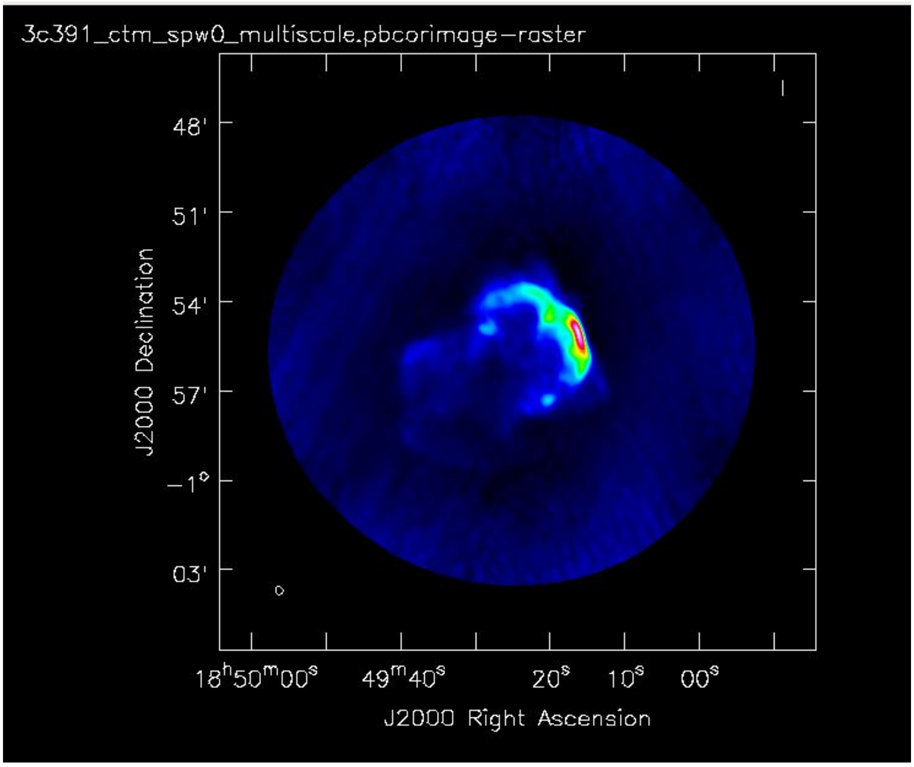


Figure 4.7: Supernova 3C391 Mask

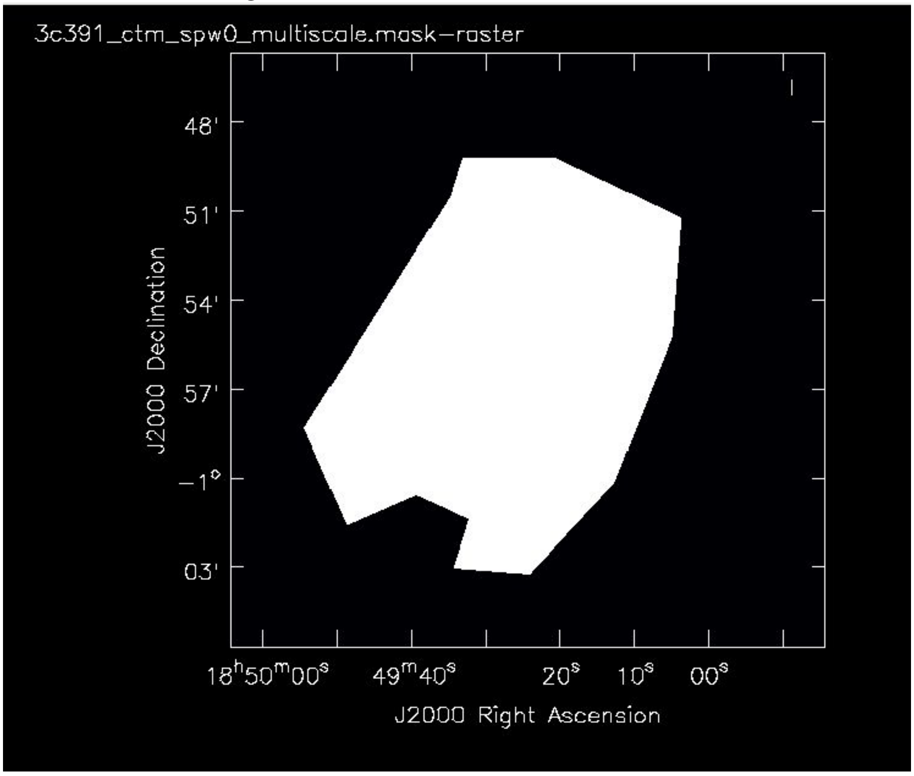
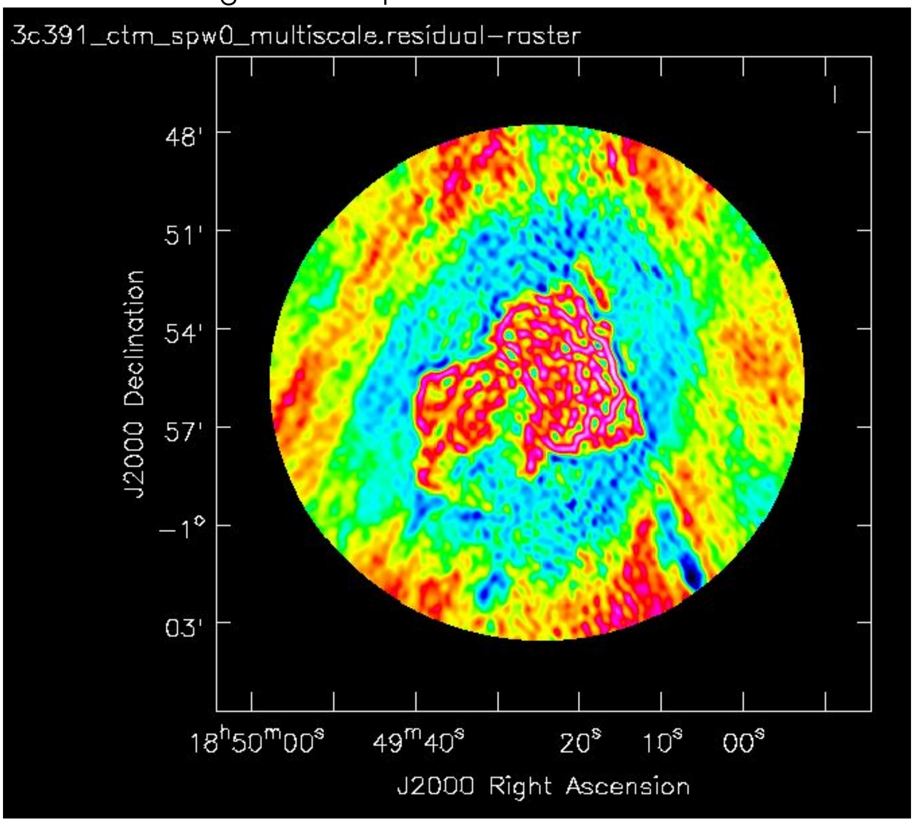


Figure 4.8: Supernova 3C391 Residual



5. Fitting the lightcurve of GW170817 using MCMC

5.1 GW170817

In section 1.4, we discussed the GW170817 event. This event led to phenomenal discoveries and verifications with the help of multi-messenger astronomy.
This event was observed by many different radio telescopes in many different frequencies.
The smooth broken power-law model given by

$$F(t,\nu) = 2^{\frac{1}{s}} \left(\frac{\nu}{3GHz}\right)^{\beta} F_p \left[\left(\frac{t}{t_p}\right)^{-s\alpha_1} + \left(\frac{t}{t_p}\right)^{-s\alpha_2}\right]^{-\frac{1}{s}}$$

fits very well to this lightcurve

5.2 Markov Chain Monte Carlo (MCMC)

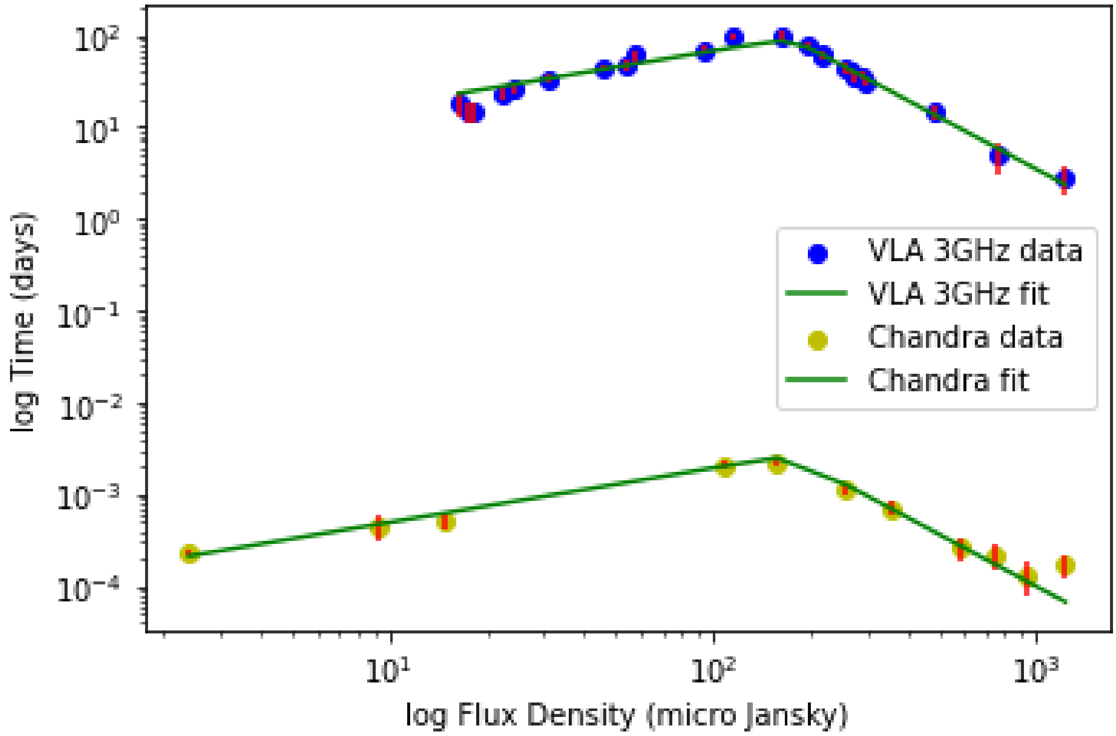
MCMC is a method for fitting models to data. To be exact, it is a parameter space exploration tool.
If the model has k parameters, we start with a set of k-dimensional vectors called walkers, which explore the parameter space. At each step, a walker gives us a model, which is compared with the data using some cost function. We then slightly change the walkers in such a way that eventually all walkers move towards the region of least cost between the model and the data.

5.3 using MCMC to fit GW170817

I used the Markov Chain Monte Carlo method to fit the smooth broken power-law model to the GW170817 related data from the VLA and Chandra telescopes.

5.4 Results

Figure 5.1: GW170817 Lightcurve calculated using Markov Chain Monte Carlo. Exact parameters mentioned in Section 5.4.



The model parameters I got using MCMC are

- $\beta = -0.575$
- $F_p = 85.622$
- $t_p = 174.096$
- $s = 7.609$
- $\alpha_1 = 0.592$
- $\alpha_2 = -1.884$

Notice that the β value obtained here is very close to the value we used in Section 1.4 (-0.584)

6. Fast Radio Bursts

Fast Radio Bursts are transient radio pulses that last from a fraction of a second to a few seconds.The cause of such bursts are not completely understood. We have observed FRBs from inside the Milky Way as well as from Extra-Galactic sources.

6.1 Dispersion

The component frequencies of each burst are delayed by different amounts of time depending on the wavelength. This delay is described by a value referred to as a dispersion measure (DM)
The DM increases as the amount of free electrons the light passes through increases.

$$DM = \int_0^d n_e dl$$

DM is equal to the number density of electrons n_e integrated along the path traveled by the photon from the source to the Earth
Hence, the DM of an FRB acts as a proxy for distance to the FRB source.

6.2 De-Dispersion

De-dispersion is a technique used in Radio Astronomy to compensate for the smearing effect of interstellar dispersion.
We know that the delay caused by dispersion at the frequency ν is given by

$$t_{delay} \propto DM\nu^{-2}$$

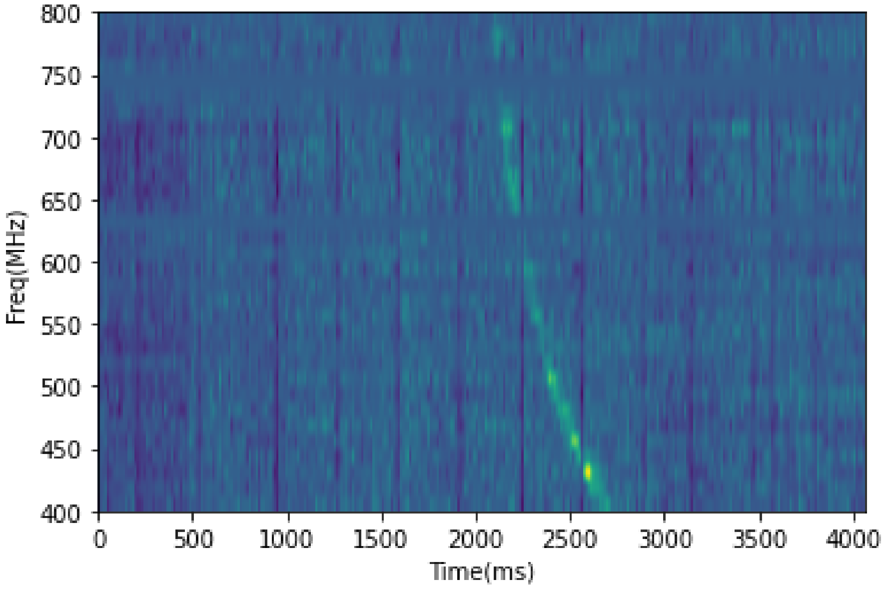
Here, the proportionality constant is 4.148808 x 10^3 if ν is in MHz, t_{delay} in in seconds and DM is in parsec cm^{-3}

6.3 Calculating DM

One way top calculate the Dispersion Measure for an FRB is to try out De-Dispersing the signal using different DM values, and seeing which one gives the highest Signal-to-Noise ratio(SNR)
I did the same for the data from a pulsar.

6.4 The Pulsar

Figure 6.1: Raw signal (Dynamic Spectrum). Notice that the signal comes at different times at different frequencies, leading to a very low SNR if we naively create a time series.



I tried DMs in the range 0 to 50 with steps of size 0.01 and got the following results:
Best DM = 30.02 parsec cm^{-3}
Peak SNR (on De-dispersing using DM = 30.02) = 12.74621

Figure 6.2: Raw signal (Time series)

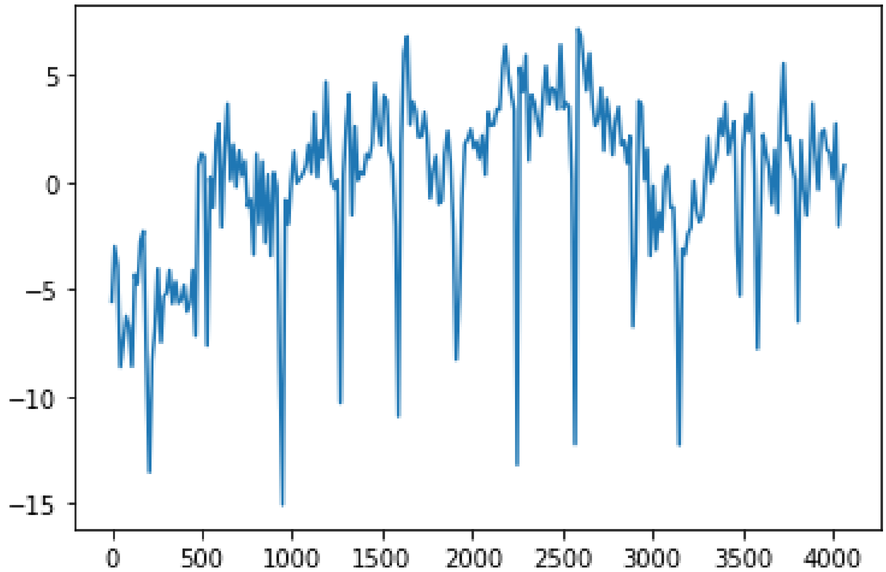


Figure 6.3: De-dispersed signal (Dynamic Spectrum). Once the signal peaks at all the frequencies are lined up, we have a better SNR.

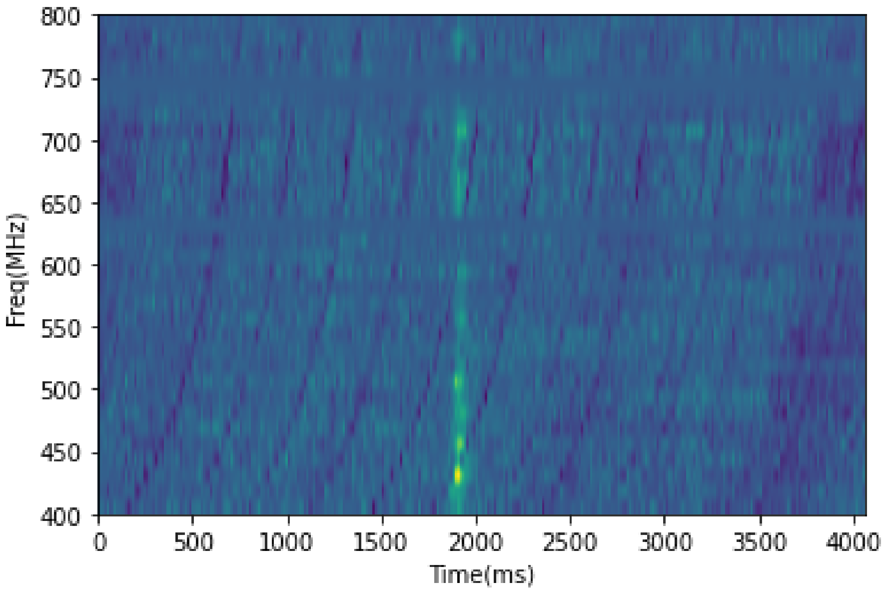


Figure 6.4: De-dispersed signal (Time series)

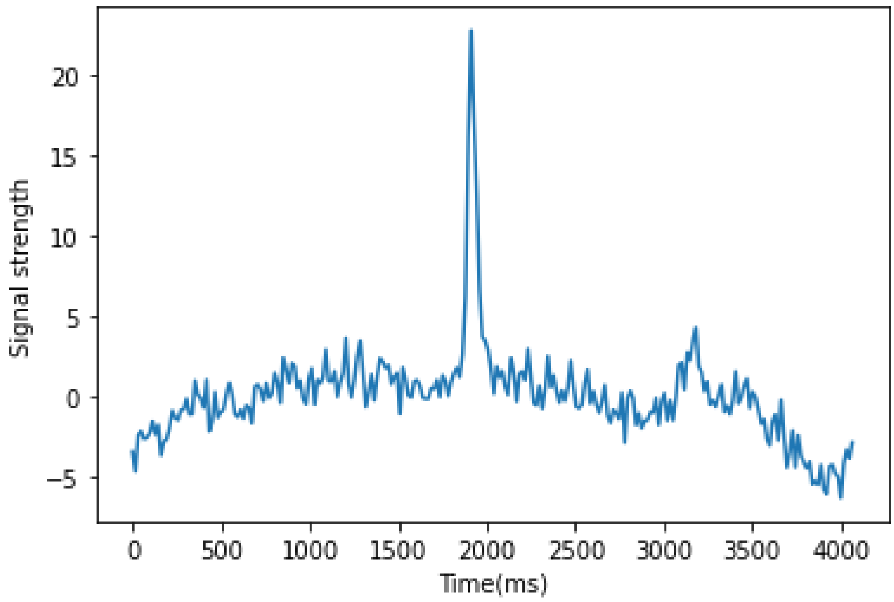
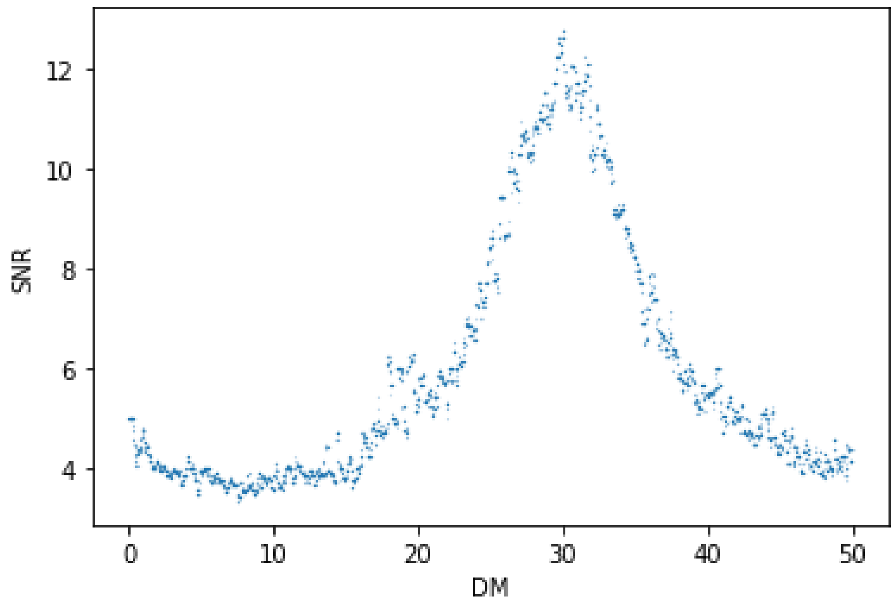


Figure 6.5: A plot of SNR vs DM. We can see that as the peak starts to line up, the SNR increases



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Books

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Articles

Python for Astronomers, MCMC Tutorial

Data

CIRADA image cutout web service
GW170817 data from Caltech

Code

Code by me (Aadyot Bhardwaj) on GitHub