The neutral hydrogen environment in the star forming region

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Abstract

G70 has always been an enigmatic nebula and has stimulated astronomers to propose the most fantastic and exotic explanations. To characterize the interstellar medium (ISM) around G70, we have obtained $H\ I$ data from the Westerbork radio telescope (WSRT). In this framework, the scope of this study has been to reduce the data, derive key parameters (optical depth, spin temperature and column density) of the surrounding ISM, whose interaction with the source is used to improve our understanding of the object. We reduced the radio data using the dedicated package ‘Astronomical Image Processing System (AIPS)’ in a standard fashion. After the uv-fits files were read into AIPS, we applied a bandpass calibration using calibrator sources 3C147 and 3C286 because the observation was initially well calibrated in both phase and amplitude. The final steps consisted in imaging and self-calibration. Nevertheless, undesired negative parts (black areas) are still left around the source on our image, which are suspected to occur from a difference between the theoretical and the real dirty beam. Subsequently, desired parameters have been deduced and the spectra obtained. Unfortunately, the beam of the $H\ I$ data was too large to be useful for a more detailed analysis of the morphology and dynamics of the cold neutral ISM around G70. Unlike our expectations, no relationship has been found between the surrounding ISM and the source.
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Introduction

In between the stars one can be found a very tenuous mixture of gas (~ 99% by mass), dust and energetic particles, referred to as the Interstellar Medium (ISM). It is of high interest, because gravitational instabilities in its cold population can lead to star formation. On the other hand, the ISM is replenished through material from planetary nebulae, stellar winds and supernova explosions. As hot newly-formed stars usually tend to heat and ionize the neighboring gas clouds, the ISM can then be found in both neutral and ionized states. In general, it is classified into 3 main categories depending on temperatures and densities: 1) cold neutral (\(T \sim 100\ \text{K}, n \sim 0.1\text{ to } 1\ \text{cm}^{-3}\)), 2) warm ionized (\(T \sim 10^4\ \text{K}, n \sim 10\text{ to } 10^4\ \text{cm}^{-3}\)) and 3) hot ionized (\(T \sim 10^6\ \text{K}, n \sim 10^{-3}\ \text{cm}^{-3}\)).

The main constituents of the ISM are by far atomic (\(H\ I\)) and molecular hydrogen (\(H_2\)). \(H\ I\) can be easily traced by its 21 cm line in the radio domain of the electromagnetic spectrum. However, molecular hydrogen being a symmetrical molecule, has no dipole moment and hence does not radiate at easily observable frequencies. Molecular gas is traced mainly by CO emission.

Rather than using space-based probes, one efficient method to investigate the ISM is by spectral analysis, carried out through ground-based telescopes. Elements are identified by their characteristic spectra. The spectra also show the dynamics of the emitting or absorbing gas by how far the lines have shifted with respect to their rest frequency.

Characterization of the ISM is a key step to have a comprehensive understanding of stars’ lifecycle, and answer pertinent questions on the structure and the evolution of our galaxy and the universe.

In this report, we describe the ISM around the galactic object G70.7+1.2. It is a peculiar object, whose structure has stimulated a noticeable amount of research, as its nature is still unknown. It shows an outflow of some kind. The work consists of the reduction of a calibrated uv-dataset of Westerbork Radio Telescope (WSRT) neutral hydrogen (\(H\ I\)) observations of G70 into an \(H\ I\) data cube. This is done by the ‘Astronomical Image Processing System’ (AIPS) software package. From this cube, we were able to reproduce the relevant \(H\ I\) spectral line, channel maps and position-velocity maps. We have also retrieved some useful parameters: \(H\ I\) column density, spin temperature and optical depth. Finally, we provided some conclusions and suggestions for possible future work.
1. **Basic knowledge in radio astronomy**

Many astronomical objects in the universe emit electromagnetic waves in the radio range. Radio waves with \(10 \, m \leq \lambda \leq 1 \, cm\) (Kraus, 1966), the so-called 'radio window', are relatively free from distortion effects caused by the Earth's atmosphere. At very short wavelengths (gamma, X-ray, UV, infrared), photons are mostly absorbed by atmospheric composition, and in the opposite way, very low energy-photons \((\lambda > 10 \, m)\), are bounced back by our ionosphere.

The atmosphere and many instruments from the ground such as mobile phones also emit electromagnetic waves in the radio range which might disrupt radio-observations. Due to the fact that dust particles are much smaller than the radio wavelength, absorption or scattering effects are considered negligible. This leads to an advantage compared to, for example, optical observations. Furthermore, many of the radio sources can be observed both day and night, without any strong disturbances from the Sun. Like most stars, the Sun is a weak emitter at radio wavelengths. Even the coolest stars emit mostly in the visible or the infrared.

Radio waves are emitted either by thermal or non-thermal mechanisms:

1) Thermal emission is strongly related to the temperature. It includes blackbody radiation, free-free emission (thermal bremsstrahlung) and line emission. The higher the temperature, the faster the molecules are vibrating or moving. Whenever electrons move in a curved path or are accelerated, electromagnetic waves are produced.

2) Non-thermal emission is produced by mechanisms unrelated to the temperature of the object. For instance, non-thermal emission can be generated by electrons spiraling around magnetic field lines (synchrotron radiation) and by the inverse compton effect.

**Blackbody radiation**

A blackbody is an object that absorbs all radiations falling upon it. When its temperature is higher than absolute zero it will emit some energy at all wavelengths. It is both the perfect absorber and emitter. More precisely, a thermal source is in thermal equilibrium with its surroundings, but a blackbody emits thermal radiation that is also in equilibrium with itself.
The blackbody spectrum can be described by *Planck's Law*:

\[
B_v = \frac{2h\nu^3}{c^2} \left( \frac{1}{e^{h\nu/kT} - 1} \right)
\]  

(1)

\(B = \text{brightness (W m}^{-2} \text{ Hz}^{-1} \text{ Sr}^{-1})\)

\(v = \text{frequency (Hz)}\)

\(c = \text{speed of light} = 2.99792458 \times 10^5 \text{ km s}^{-1}\)

\(h = \text{Planck's constant} = 6.63 \times 10^{-34} \text{ J s}\)

\(k = \text{Boltzmann constant} = 1.3806503 \times 10^{-23} \text{ J K}^{-1}\)

\(T = \text{temperature (K)}\)

In case of radio waves with very long wavelengths \((\nu/kT \ll 1)\), we can replace the exponential term in Planck's law by Taylor-series approximation,

\[
\exp \left( \frac{\nu}{kT} \right) - 1 = 1 + \frac{\nu}{kT} + \cdots - 1 \approx \frac{\nu}{kT}
\]  

(2)

yields the so-called Rayleigh-Jeans approximation,

\[
B(T) \approx \frac{2h\nu^3}{c^2} \frac{kT}{\nu} = \frac{2kT \nu^2}{c^2} = \frac{2kT}{\lambda^2}
\]  

(3)

In general, we can write

\[
B_\lambda \propto \lambda^n
\]  

(4)

\(n = \text{spectral index (it is simply the slope between } \log B_\lambda \text{ and } \log \lambda)\)

The spectral index for the thermal radiation from a blackbody is then equal to \(-2\) and for most non-thermal sources the spectral index lies between about +0.3 and +1.3 (Kraus, 1966).

By simple mathematics, the wavelength at which the brightness is maximum \((\lambda_{\text{max}})\) can be calculated by,

\[
\frac{\partial B_\lambda}{\partial \lambda} = 0
\]  

(5)
\[ \lambda_{\text{max}} = \frac{2.898 \times 10^{-3}}{T} \]  

(6)

This equation is called *Wien's law*. The blackbody with a higher temperature will have a shorter wavelength at the peak of the radiation's brightness.

With more advanced calculations we can find the total brightness emitted by the blackbody at all wavelengths which is the area under the curve,

\[ B(T) = \int B_v(T) \, dv = \frac{\sigma T^4}{\pi} \]

(7)

This is the *Stefan-Boltzmann's law* and below is the Stefan-Boltzmann constant,

\[ \sigma = \frac{2\pi^5 k^4}{15c^2 h^3} \approx 5.67 \times 10^{-5} \text{erg cm}^{-2} \text{s K}^4 \text{Sr} \]

(8)

so the hotter the blackbody, the more energy it emits, very strongly dependent on temperature!

Note here that brightness (or specific intensity, \( I_v \), at the emitting surface of the source) and flux density are not the same.

Brightness is measured in the unit of \( W \text{ m}^{-2} \text{ Hz}^{-1} \text{ Sr}^{-1} \),

\[ B_v = \frac{dP}{\cos \theta d\sigma dv d\Omega} \]

(9)

This quantity does not rely on the distance between the source and the observer but flux density does (inverse-square law),

\[ S_v \propto r^{-2} \]

(10)

Flux density is in the unit of \( W \text{ m}^{-2} \text{ Hz}^{-1} \), or in Jansky (1 Jy = \( 10^{-26} \text{ W m}^{-2} \text{ Hz}^{-1} \)),

\[ S_v = \int B_v(\theta, \phi) \cos \theta d\Omega \]

(11)
Free-free emission (thermal bremsstrahlung)

A free electron is accelerated towards a positively-charged ion and the electron emits a photon in the form of radio radiation. This electron remains free both before and after the emission.

In the region around an O or a B star, the ionized hydrogen atoms form a so-called $H\ II$ region. The free electrons from these $H\ II$ atoms move independently in this region and they often interact with the ions, emitting the so-called free-free radiation. Due to the low kinetic energies of the electrons, $H\ II$ regions are a source of radio emission.

![Bremsstrahlung emission](image)

*Figure 1: Bremsstrahlung emission*

Line emission

This involves changing energy states of electrons either from upper to lower state (emission line) or lower to upper state (absorption line). In case of a neutral hydrogen atom ($H\ I$), normally an electron’s spin is not in the same orientation as the proton’s spin and this situation occupies a low energy (ground state) or low magnetic moment. But it can happen that the directions of both spins become parallel with a slightly higher energy or higher magnetic moment. This takes place when the atom collides with another atom or electron. After some time the atom returns spontaneously to the ground state and it loses the energy by emitting a 21.11 cm wavelength photon. This is called changing in hyperfine states of the $H$ atom (more details in the chapter of ‘$H\ I$ 21 cm line’).
Spectral lines are never completely sharp but they are always broadened! There are several reasons for this e.g.:

1. **Natural broadening:** From the Heisenberg uncertainty principle, this broadening depends on the lifetime of an energy level.

   \[
   \Delta E = \frac{1}{2\pi} \frac{\hbar}{\Delta t}
   \]

   \(12\)

2. **Doppler (thermal) broadening:** This broadening relates to the temperature. Red shifts happen when particles move away from the observer and in the opposite way, blue shifts occur when particles move toward the observer. The various velocities of particles lead to a broadening of the spectral line.

   \[
   \nu = c \left(1 - \frac{\nu}{\nu_0}\right)
   \]

   \(13\)

   \(c = \text{speed of light} = 2.99792458 \times 10^5 \text{ km/s}\)

   \(\nu = \text{frequency}\)

   \(\nu_0 = \text{rest frequency}\)

3. **Collisional broadening:** Density influences this kind of broadening.

4. **Zeeman effect:** As an electron has their own intrinsic magnetic dipole moment, so its energy can be separated into many sub-levels when it stays in the magnetic field.
**Synchrotron radiation**

When an extremely high energy charged particle (mostly electrons) moves in a magnetic field, the particle is forced to move spirally around the magnetic lines of force. Once the particle is accelerated it radiates electromagnetic radio waves.

![Figure 3: Synchrotron radiation](image)

**Inverse compton effect**

This occurs when an electron has energy comparable to that of a photon. As the electron collides with the photon the electron transfers its energy to the photon causing the emission of a photon with higher energy.

![Figure 4: Inverse compton scattering](image)
Radio maps

Spectral line graphs

A spectral line graph is usually a plot of intensity as a function of velocity. The velocity is determined by Doppler relationship, from this we know how fast gas is moving toward and away from us. For example, the spectral line of $H\,I$ ‘at rest’ is $\sim 1420$ MHz. But due to the rotation of the Galaxy, relative motions of gas with respect to Earth will appear and their spectra lines will hence be Doppler shifted. Because the Earth’s motion around the Sun varies with time, we count observed velocities for heliocentric velocities. The Sun will have a relative velocity with respect to the average velocities of the stars in its neighbor. This velocity is called ‘local standard at rest’ (LSR).

Channel maps

Objects in space emit radiations in a wide range of frequencies. Most spectral line observations focus only on a particular narrow range of frequencies. The bandwidth is divided into frequency channels. Each channel map shows a gas moving at only one narrow velocity range. The chosen area for a channel map is represented by two spatial coordinate: right ascension and declination. The intensity of the source is specified either by the different colors on the image or by contours on a contour map.

Average image

When we average all channel maps together, an ‘average image’ will be created so the average image does not have the velocity information. It is, however, a good way to see the morphology and how gas extends within the source.

Position-velocity diagrams (PV-diagram)

A ‘PV-diagram’ displays how gas travelling at various velocities is dispersed along a narrow line-of-sight range. And again colors or contours indicate the intensity of the emission distributed along the line-of-sight. The PV-diagram provides more details to understand the relative motions of gas in space such as rotations and expanding shells.

Data cube

A data cube is simply a three dimensional pile of all channel maps. In my case, two axes are spatial coordinates and the third one is velocity. Rather than seeing all channel maps sequentially by the channel map movie, they are all able to be seen at once. And intensity is indicated by a numerical value.
2. **Radio interferometry and aperture synthesis telescopes**

In order to study radio source with a single dish radio telescope likewise in optical observations, there is a main limitation of achieving the high resolution image: the diameter of the telescope must be many times larger than the wavelength of the observed object,

\[
\theta = 206265 \times 1.22 \frac{\lambda}{D}
\]

(14)

\(\theta\) = resolution (arcsec)
\(D\) = aperture of the telescope
\(\lambda\) = wavelength

With the interferometry theory, it is feasible to 'synthesize' a filled-dish aperture with many small telescopes and acquire almost the same resolution as the enormous one. This technique is called 'aperture synthesis'. To simply understand the fundamental approach of 'radio interferometry and aperture synthesis', first consider only two antennas (imagine the double-slits experiment):

As we can see in the Figure 5\(^1\), the covered distance from the source to the antennas of these two radio signals are not the same. The signal must travel \(\delta s\) more to the left antenna and the calculated time delay for this,

\[
\delta t = \frac{\delta s}{c} = \frac{R \sin \theta}{c}
\]

(15)

\(c = \text{speed of radio wave} \approx 2.99792458 \times 10^8 \text{ m s}^{-1}\)
\(R = \text{baseline}\)
\(\theta = \text{angle between the normal baseline and the source}\)

---

\(^1\) In this section, I follow the notes on 'Theory of interferometry and aperture synthesis' by Migenes and Andernach (http://www.astro.ugto.mx/cursos/RadioAstronomy/Radioastronomy-4.pdf)
Let $A$ be a received signal from antenna 1,

$$A = A_1 e^{i\omega t}$$  \hspace{1cm} (16)

and $B$ from antenna 2,

$$B = B_1 e^{i(\omega t - k\delta s)}$$  \hspace{1cm} (17)

$\omega$ = angular frequency
$k$ = wavenumber

The output of the multiplier ($M_{12}$) is then,

$$M_{12} = AB^* = A_1 B_1^* e^{ik\delta s}$$

$$= A_1 B_1^* e^{ikR\sin\theta}$$  \hspace{1cm} (18)

Now, to attain the combined signal from the integrator ($V_{12}$): we have to first consider a source brightness distribution $B_S(x)$ for only 1 dimension where $x$ is the displacement angle from the field center at $\theta$,

$$V_{12} = \int B_S(x)e^{ikR\sin(\theta+x)}dx$$  \hspace{1cm} (19)
For small value of $x$, $\sin(\theta + x) = \sin \theta + x \cos \theta$

$$V_{12} = e^{ikR\sin \theta} \int B_S(x)e^{ikR\cos \theta} \, dx$$

(20)

Since the signals from both antennas have to be constructively interfered, the adding of this delay must be taken into account, which is commonly redeemed electronically. For the case of many baselines, this compensation is determined only for very short time interval because the distances between the source and each telescope change all the time.

Hence the term $e^{ikR\sin \theta}$ which is the result of time delay can be removed from our discussion. And we can also write,

$$kR \cos \theta = 2\pi u$$

(21)

$u =$ spatial frequency

The correlator output becomes,

$$V_{12} = \int B_S(x)e^{2\pi iux} \, dx$$

(22)

If we extend this into 2 dimensions in a similar manner,

$$V_{12} = \int B_S(x, y)e^{2\pi i(ux + vy)} \, dxdy$$

(23)

This is a so-called ‘Fourier transform’. If we do an inversion of this transform,

$$B_S(x, y) = \int V_{12}(u, v)e^{-2\pi i(ux + vy)} \, dudv$$

(24)

it will yield a source brightness distribution. So, the measured signals on a uv-plane (Fourier transform or frequency domain plane) can be transformed to the source brightness distribution (spatial domain) by the inversed Fourier transform.

Each $(u, v)$ coordinate, which is known as a visibility $V(u, v)$, agrees to one pair of telescopes. The more the uv-plane is filled, the better images are produced. Therefore a large number of different baselines is required (multi-slits interferometer). The word 'filled' is used here because, of course, there are still a lot of gaps in between
which leads to many sidelobes. And also, the technique of interferometer causes also the so-called ‘zero-spacing problem’. Since the baseline is always finite, the flux density at the position (0, 0) on the uv-plane can never be measured. For a point source, it is easy to estimate this value from the smallest \(u\) and \(v\) since the visibility as a function of baseline length is constant. But for an extended source, it is difficult to do this even with extrapolation.

In most cases, to save the cost of radio antennas the trick of using the rotation of the Earth is introduced: in the course of a 12 hour-observation, a full circle or ellipse will be created and simulate the immense single-dish radio telescope in the end.

Three interesting properties of interferometry and aperture synthesis radio telescopes are listed here:

1. Resolution : It is determined by the longest baseline (the diameter of the single-dish telescope).

2. Sensitivity : The number of baselines (total collective area) influences this parameter. \(N\) telescopes give \(\frac{N(N-1)}{2}\) baselines.

3. Field-of-view : This one does not depend at all on how much the \(UV\)-plane is filled nor the length of the longest baseline but relies directly on the size of each single dish telescope itself. The larger the aperture, the narrower is the beam.

An example of the telescope with this technique is Westerbork Radio Telescope.

**Westerbork Radio Telescope**

The WSRT comprises 14 dish-shaped telescopes which are arranged on a 2.7 km East-West line. Each telescope has a diameter of 25 meters. Four out of these fourteen antennas are movable, the rest have their own fixed positions. The technique of interferometry and aperture synthesis is being employed here. The antenna dish collects the radiation falling on the whole surface area and then focus to the frontend (radio receiver). The radio signals from all 14 antennas will be delivered together through the coaxial cables to the control room and will be operated to get a high resolution image there. The WSRT observes the sky usually in the course of 12 hours and with the Earth’s rotation, it creates a full circle reproducing a huge single-dish telescope.

WSRT is organized by the 'Netherlands Foundation for Research in Astronomy, NFRA'. The location is near the village of Hooghalen in the area of the 'Voormalig Kamp Westerbork', Netherlands.
3. The H I 21 cm line

Hydrogen is the most important element to study in the ISM because of its large amount. As the H$_2$ molecule is symmetric, there is no permanent dipole moment and thus no noticeable radio spectral line emission. But neutral hydrogen (H I), can be detected in the 21cm (1420.405752 MHz) H I transition or hyperfine line.

The hydrogen atom consists of 1 proton and 1 electron. The electron has a spin \( S \) of 1/2 and the proton has a nuclear spin (or nucleus total angular momentum \( I \)) of 1/2, hence the total spin (\( F \)) of the atom can be either 0 or 1. This leads to the ground level splitting corresponding to the energy of \( 5.87 \times 10^{-6} \) eV or \( 9.41161 \times 10^{-25} \) J.

The frequency of this hyperfine line is calculated by

\[
v = \frac{8}{3} g_I \left( \frac{m_e}{m_p} \right) \alpha^2 (Rc)
\]

\( g_I \) = nuclear g-factor (ratio of magnetic moment and angular momentum) 
\approx 5.58569 \ (for \ proton)

\( m_e \) = electron mass \approx 9.10938188 \times 10^{-31} \) kg

\( m_p \) = proton mass \approx 1.67262158 \times 10^{-27} \) kg

\( \alpha \) = fine – structure constant = \( \frac{e^2}{(hc)} \approx \frac{1}{137.036} \)

\( R \) = Rydberg constant = \( \frac{m_e e^4}{8\pi^2 \hbar^3 c} \approx 1.09737316 \times 10^7 \) m$^{-1}$

\( c \) = speed of light \approx 2.99792458 \times 10^8 \) m$^{-1}$

So,

\[
v = \frac{8}{3} (5.58569) \left( \frac{1}{1836.15} \right) \left( \frac{1}{137.036} \right)^2 (3.28984 \times 10^{15} Hz)
\]

\approx 1420.405751 \) MHz

---

\(^2\) In this chapter, I follow the notes on ‘The Hi line’ by Condon and Ransom (http://www.cv.nrao.edu/course/astr534/HILine.html) and also the notes on ‘ISM’ by E.F. van Dishoeck (http://www.strw.leidenuniv.nl/~dave/ISM)
Then the transition energy,

\[ \Delta E = h\nu = (6.626 \times 10^{-34}) (1420.405751 \times 10^6) \]
\[ \approx 9.41161 \times 10^{-25} \text{ J} \]
\[ \approx 5.87 \times 10^{-6} \text{ eV} \]  

(27)

\( h = \text{Planck’s constant} \approx 6.626 \times 10^{-34} \text{ m}^2 \text{ kg s}^{-1} \)

With Boltzmann’s distribution, the mean relative population of electrons in upper and lower state can be obtained by

\[ \frac{n_U}{n_L} = \frac{g_U}{g_L} \exp \left( -\frac{h\nu}{kT_S} \right) \]

(28)

\( n_U = \text{number of electrons in upper state} \)
\( n_L = \text{number of electrons in lower state} \)
\( T_S = \text{spin temperature (K)} \)
\( k = \text{Boltzmann constant} \approx 1.3806503 \times 10^{-23} \text{ m}^2 \text{ kg s}^{-2} \text{ K}^{-1} \)

The statistical weight of upper state \( (g_U) \) and lower state \( (g_L) \) are,

\[ g_U = 2F + 1 = 2(1) + 1 = 3 \]
\[ g_L = 2F + 1 = 2(0) + 1 = 1 \]  

(29)

In practical, \( kT_S >> h\nu \)

\[ \frac{n_U}{n_L} = \left( \frac{3}{1} \right) (1) = 3 \]  

(30)

The temperature that accounts for the distribution of the atoms in upper and lower states is called spin or excitation temperature \( (T_S) \). As the hyperfine transition is caused mostly by collisions, normally \( T_S = T_{kin} \) (kinetic temperature).

The Einstein spontaneous emission coefficient \( (A_{UL}) \) of this transition (in the unit of CGS) is defined as,

\[ A_{UL} \approx \frac{64\pi^4}{3hc^3} v_{UL}^3 |\mu_{UL}|^2 \]  

(31)
$|\mu_{UL}|$ is the Bohr magneton (a unit of magnetic moment)

$$\mu_{UL} = \frac{e\hbar}{2m_e c}$$

$$\approx 9.27 \times 10^{-21} \text{ erg Gauss}^{-1} \quad (32)$$

Hence,

$$A_{UL} \approx \frac{64\pi^4(1420.405751 \times 10^6 \text{Hz})^3(9.27 \times 10^{-21} \text{erg Gauss}^{-1})^2}{3(6.626 \times 10^{-27} \text{erg s})^3(3 \times 10^{10} \text{ cm s}^{-1})^3}$$

$$\approx 2.85 \times 10^{-15} \text{ s}^{-1} \quad (33)$$

The radiative half-life is then

$$\tau_{1/2} = A_{UL}^{-1} = \frac{1}{2.85 \times 10^{-15}} \approx 11 \text{ million years} \quad (34)$$

Although $\tau_{1/2}$ is extremely long, hydrogen is the most abundant element in the universe, present almost everywhere. So we are able to observe its 21 cm spectral line in recognizable strength.

**Optical depth and column density**

The relationship between the optical depth and the column density of neutral hydrogen atoms is defined by

$$N(H I) = 1.818 \times 10^{18}(\tau_v T_S) \Delta \nu \quad (35)$$

$N(H I) =$ the $H I$ column density in cm$^{-2}$

$\Delta \nu =$ the velocity width of the observed emission in km s$^{-1}$
$T_S$ has an inverse link with $\tau_u$ so warm $HI$ regions (tenuous clouds) cannot be measured in absorption or we can say that higher $T_S$ means more electrons in the upper state (more non-absorbing atoms).

In an optically thin case, brightness temperature $(T_B) \approx \tau_u T_S$,

$$N(HI) = 1.818 \times 10^{18} \int T_B d\nu$$

If there is a continuum source behind an $HI$ cloud, we will be able to solve the optical depth from this equation (which is the absorption component in $HI$ spectra):

$$T_{annulus} - T_{source} = T_{cont} (1 - e^{-\tau})$$

(37)

$T_{annulus}$ = the brightness temperature of the emission in the annulus around the continuum source.

$T_{source}$ = the brightness temperature of the beam centered on the continuum subtracted source.

$T_{cont}$ = the brightness temperature of the continuum source.

Once we know the optical depth, we can determine the spin temperature by,

$$T_{source} = T_{cont} (e^{-\tau} - 1) + T_S (1 - e^{-\tau})$$

(38)

Note that the flux density is associated with the brightness temperature by Rayleigh-Jeans law in Chapter 1 of the present report.
4. **G70.7+1.2**

G70 is a compact galactic radio and optical nebula with a shell-like radio structure (Minkowski, 1948; Reich et al., 1984; Green, 1985). It is enclosed by a lumpy pattern of molecular clouds or globules (Bally et al., 1989). This source displays a very particular physical nature. Many explanations have been put forward to explain this object including a young supernova remnant, an ionized hydrogen region, a nova shell and a young stellar object outflow (Reich et al., 1985; Green, 1986; de Muizon et al., 1988; Bally et al., 1989). However, none of these theories managed to give a reason for low expansion velocities with the strong non-thermal radio emission of this source.

In 1984 and 1985, Reich et al. proposed that this source has a supernova origin due to its characteristic non-thermal high frequency spectrum, its shell-like morphology and its polarization. Its high surface brightness and compact size led to the conclusion of a very young supernova remnant. This conclusion was not supported by Green (1986) as: no significant polarization was found, the shape of this source was not really circular and the infrared spectrum was similar to that of compact H II regions. Therefore he suggested that G70 should be a thermal emission nebula rather than a very young supernova remnant.

Two years later, a study in various wavelengths, radio, infrared, optical and X-ray, was carried out by de Muizon et al. They confirmed the presence of shocked gas in a very dense environment. The radio observations revealed non-thermal synchrotron emission with a weak polarization of 1.6% at 6 cm. The non-detection of recombination lines at 2 cm, did not support the presence of an H II region. They proposed that G70 might be a nova remnant.

In 1986 Bally et al. posed that G70 might be powered by energetic outflows from a young stellar object. Nevertheless, the high level of radio continuum and polarization could not be explained by this model. With CO observations, they estimated the distance to the object to $4.5 \pm 1$ kpc.

A few years after, Kulkarni et al. (1992) detected a Be-star (B-type star with one or more emission lines of hydrogen in its spectrum) near the bow shock creating a reflection nebula. They proposed that a Be-star/pulsar binary system could explain the observations. The released energetic particles from a Be-star through a dense material create a bow shock resulting in the optical emission. The non-thermal radio emission comes from shocked supersonic pulsar winds.

Recently, G70 was suggested to be the combination of two unrelated objects, A Be-star and X-ray emitting B-star/pulsar binary, interacting with a molecular cloud (Cameron and Kulkarni, 2007). This model was strengthened by the discovery of an X-ray source with a NIR counterpart case (which should be the B-star) in the centre of G70.
5. Data reduction and related works

We reduced the data with a software package called 'Astronomical Image Processing System' (AIPS). It is a software package for interactive calibration, image construction, image display, data analysis, plotting and a diversity of necessary tasks on Astronomical data obtained by radio interferometry. It performs analysis of astronomical images made from data by using Fourier synthesis methods.

Many subprograms or applications can be run as a series or in parallel after original setup of parameters. Those are defined as separate ‘tasks’ that are to be executed in time depending on the user’s need. The length of those tasks can differ significantly and use various amount of CPU memory. Work can be done in parallel that other tasks are running. ‘Adverbs’ are responsible for passing parameters to ‘tasks’. They can be real numbers, character strings, scalars or arrays.

There are four main steps for data reduction:

1) Calibration
2) Construction of a so-called ‘dirty image’
3) Deconvolution
4) Self-calibration

The observation dates for our data were on the 17th January 1987 and 17th April 1987 by Westerbork Radio Telescope (WSRT). The spatial resolution of the map is \(33.51'' \times 15.64''\) and the velocity width is 2.07280 km/s.

1) Calibration

The first procedure of data processing in aperture synthesis is the calibration. We do the calibration to reduce errors that might be caused by atmospheric or instrumental instabilities.

Since our data has already been initially calibrated both in phase and amplitude, we started the procedure with a ‘bandpass’ calibration which is the calibration of the response of the system across all channels within which data have been recorded. The antenna bandpass functions are determined from calibrator data.
Bandpass with the calibrators 3C147 and 3C286

*task ’BPASS’*: To create a bandpass (BP) table that contains the BP response functions of the antennas. The bandpass solutions, determined by ’BPASS’, are written to a BP table extension of the input uv-dataset. To run this task we specified:

1) The input uv-dataset which are calibrator sources 3C147 and 3C286.
2) The antenna to determine the bandpass. All antennas for our case.
3) The time interval over which to average the uv-data before solving for the bandpass. We chose to average the uv-data for the whole time range.

*task ’POSSM’*: To plot the BP spectra generated by task ’BPASS’. For this task, we determined:

1) The input uv-file name containing the BP table.
2) We set the adverb ’Doband’ = 1 meaning that we are applying the created BP table to the data

The results are shown in Figure 6 and 7. The amplitude part of these figures (lower plot) tells about the sensitivity of each channel. We are then able to see that channel 1 and channels 120 to 128 are not sensitive during these observation times. The phase part (upper plot) tells about how good time delay is compensated by the system. Relevant correction is applied from the combination of the amplitude and phase BP plots.
Figure 6: Bandpass plotting for day 1 (all antennas averaged)
Figure 7: Bandpass plotting for day 2 (all antennas averaged)
Copy the BP tables

Next step, we copied the created BP tables from the calibrator sources 3C147 and 3C287 to the object of our interest, G70.7+1.2, for both day 1 and day 2 respectively,

*task ‘TACOP’*: To copy the table-type files. The specified output file should not exist before the execution of ‘TACOP’. Multiple files can be copied and table copies are conditioned from the specific input parameters.

Combining the two day-observation data

Before start merging our two day-data, we have to first apply the BP tables permanently to both day 1 and day 2 uv-data. This is done by *task ‘SPLIT’*.

*task ‘SPLIT’*: To apply an unapplied table to an input file. This is done by copying a table to the input file.

After this, the data are combined with *task ‘DBCON’*,

*task ‘DBCON’*: To combine two datasets taken at different times. These datasets should contain observations of roughly the same frequency.

2) Construction of a dirty image

A telescope does not actually construct the image by itself but it gains the observed information (uv-data) which is used to make the image instead.

First, the available uv-data is transformed into a form suitable for the Fourier inversion by gridding and weighting. Gridding involves an interpolation of available data points and a decision on what value to use where no data points exist. Usually, it adds the Fourier components with some form of weighting. Weighting can make a big difference in the signal-to-noise ratio on images either by altering the beam width or a sidelobe pattern. For pure uniform weighting, the beam width is very narrow but contains high sidelobe levels and for pure natural weighting, the resolution is less but
the noise levels decrease. Robust weighting lies somewhere in between uniform and natural weighting. Finally, the ‘dirty image or dirty beam’ will be obtained by applying the inversed Fourier transform to the gridded and weighted uv-data (Figure 9). Without further correction, the map still contains large sidelobes.

3) Deconvolution

However, the precise form of the dirty beam is known from the distribution of the uv-data. The ‘CLEAN deconvolution’ starts with the subtraction from the dirty map, at a peak position, of a dirty beam with suitable amplitude and centered on the source. Each subtraction the noticeable sidelobes will be less, revealing possibly weaker sources. The same process will be repeated until no sources can be distinguished from the noise. The ‘CLEAN’ map can now be reconstructed by adding all the individual components back to the CLEANED image, using an idealized point spread function (PSF) (Figure 10).

The construction of the dirty image and the deconvolution are made with task ‘IMAGR’. This task is intended to be the only imaging task in AIPS. It offers the full range of weighting, gridding and deconvolution. The output file is CC extension.

Below is an example of our input parameters for this task:

- **docalib 1**: To choose if we want to apply a self-calibration (SN table) to the uv-data or not. Since we have done already the self-calibration but not yet applied to the uv-data, input is 1.
- **doband -1**: To choose if we want to apply a bandpass (BP table) to the uv-data or not. We already applied the BP table to the uv-data (with task ‘SPLIT’) so the input is -1.
- **bchan 2**: The first channel to start imaging
- **echan 119**: The last channel to image
- **nchav 1**: The number of channels to be averaged together contributing to each output image.
**cellsize 4, 4:** Separation between pixels of a 2-dimensional image. The first element is the increment between columns (X) and the second element is the increment between rows (Y). It is calculated in the following way,

The resolution is given by:

\[ \theta = 206265 \times 1.22 \frac{\lambda}{D} \]

\( \theta \) = resolution (arcsec)
\( D \) = longest baseline = 2700 m (for WSRT)
\( \lambda \) = wavelength = 0.21 m (for H I observation which is our case)

So,

\[ \theta = 206265 \times 1.22 \frac{0.21}{2700} = 19.572258 \text{ arcsec} \]

And 1 pixel should be about 3-5 times less than the resolution,

\[ \text{cell size} \sim \frac{19.572258}{5} \sim 3.9144561 \sim 4 \text{ arcsec} \]

**imsize 2048, 2048:** Size of a 2-dimensional image to produce or process. The first element is usually called the X coordinate and measures the number of columns in the image. The second element, the Y coordinate is a measure of the number of rows in the image. It is computed in the following way,

\[ \text{field - of - view} = 206265 \times 1.22 \frac{\lambda}{D} \]

\( D \) = the diameter of each antenna = 25 m (for WSRT)
\( \lambda \) = wavelength = 0.21 m
So,

\[
\text{field-of-view} = 206265 \times 1.22^{0.21/25} = 2113.8037109 \text{ arcsec}
\]

However, it is commonly admitted to choose three times bigger than this value to include wider area since it is possible that there are sometimes very bright sources nearby,

\[
\text{image size} = \left(\frac{2113.8037109 \times 3}{4}\right) = 1585.3527832
\]

The image size must be a certain number \(2^N\), so we set the image size as 2048, 2048.

\textit{uvrange 0.2 0:} Specify minimum and maximum projected baselines to be included. The unit is in kilolambda \((k\lambda)\).

After checking with task ‘UVPLT’, we see that a few shortest baselines (~ 0 to 0.2) show a rapid increase in flux density (Figure 8), which is not supposed to happen and is considered as an error: observation of a point source such as in our case should return a flux density of nearly the same value across all baselines. Such uv-data has been excluded.

\textit{robust 0.5:} The robustness parameter is used to temper the effects of uniform weighting. The value -5 means pure uniform weighting and 5 for natural. After performing a trial-and-error, the value of 0.5 proved to be the most suitable compromise between noise and resolution.

\textit{niter 2000:} Number of clean iterations to be attempted or a number representing the maximum number of iterations in a cleaning an image process. We stopped cleaning when 2000 components have been subtracted since the peak flux density is already at the noise level.
The x-axis unit is kilo wavelength, with the $H\, I$ spectral line as reference (21cm) and kilo meaning a multiplication by a factor of 1000: 210m.
4) Self-calibration

After performing the regular calibration process, some errors might still be present. The reasons can be that the effects of the atmosphere are not the same across the sky, and also it is not possible to observe the calibrator sources all the time. That’s why we need to do ‘self-calibration’ to remove noise caused by atmospheric fluctuations.

The atmospheric perturbation affects independently each antenna, which will then generate a specific signal response. Self-calibration method uses some of the data collected by each antenna to determine more accurately its response and another part of this data is combined with signals collected by the other antennas in order to create the image.

To do self-calibration, it requires first a CLEAN model (a CLEAN image from task ‘IMAGR’) and also the visibilities that would be produced by this model (Fourier Transform). These visibilities will be different from the observed (calibrated) visibilities and one can vary the complex gains $g_i$ to minimize the differences between the model visibilities $V_{ij}^{mod}$ and the calibrated visibilities $V_{ij}^{cal}$. Schwab’s procedure is to minimize the differences in a least-square sense (Burke and Graham-Smith, 2002):

$$\delta \sum_k \sum_{i,j} w_{ij}(t_k) |V_{ij}^{cal} - g_i(t_k)g_j^*(t_k)V_{ij}^{mod}(t_k)|^2 = 0$$  \hspace{1cm} (39)

$w_{ij}(t_k) = \text{weights applied to the various interferometer pairs for each observation at } t_k$.

High signal-to-noise visibilities should receive higher weight than low signal-to-noise visibilities. The corrections to the complex gains can then be applied to the initial data set, to give a better approximation to the corrected visibilities, and the mapping procedure is then applied to the new data set.

Therefore, instead of calibrating with the already known calibrators, self-calibration uses the observations from the unknown source itself to improve the image of that unknown source.
And like every time, AIPS always make things much easier. Self-calibration is
done by task ‘CALIB’. This task determines the calibration to be applied to a uv-dataset
given a model of the source. The output data will have the corrections applied for the
input source file and the solutions that task ‘CALIB’ generates are stored in SN tables.
The SN tables will then be attached to the input file.

However, with all of these processes, there are still some negative parts (dark
areas) left around the source that we could not get rid of. This happens because the
theoretical dirty beam is probably not exactly the same as the dirty beam (Figure 11).

Continuum subtraction with AIPS

In our case, the strong non-thermal radio continuum arises from G70.7+1.2.
Normally, to detect a weak spectral line (such as 21-cm line emission and absorption) in
the presence of this strong continuum can be very difficult. The problem is related to
the ‘spectral dynamic range’ (SDR). SDR is the ratio of the weakest spectral feature that
can be detected to the continuum flux density. This can be limited by various reasons,
for example instrumental or atmospheric instabilities. With the continuum subtraction,
SDR will be determined by the peak spectral line flux density rather than the continuum
flux density. Apart from the continuum flux density, any other things that have a
constant or linear variation across frequency will be subtracted in the continuum
subtraction procedure leading to the better SDR.

The suitable task for continuum subtraction in AIPS is task ‘UVLSF’.

*task ‘UVLSF’:*  
It fits a straight line to selected
channels and subtract the fitted line
from the spectrum.

Our important inputs for this task:

docalib 1:  
To apply the SN table permanently.

Ichansel 2, 32, 1, 0, 92, 119, 1, 0:  
To select groups of channels to fit as
sets of (start, end, increment,
channel numbers). In our case, there
are two continuum regions,
channels 2 to 32 and 92 to 119,
applying to all channels.
After all the data reduction procedures, we will eventually attain the \( H I \) data cube. Nevertheless, for all plotting manners we used a software known as ‘Interactive Data Language’ (IDL). So we had to export the data cube out of AIPS by task ‘FITT’, which creates a so-called ‘fits’ file.

‘Fits’ stands for Flexible Image Transport System. It is a standard format developed in order for most astronomical software to be able to read the same data, independently of the way each of them calculates or defines parameters. All major astronomical software packages read and write ‘fits’ format data, and many have adopted ‘fits’ for exchange with other programs and facilities.

**Export and display channel maps outside AIPS**

We employed task ‘KNTR’ to produce the grey-scale plots of the images. Then the PostScript files of these maps are written by task ‘LWPLA’.

*Task ‘KNTR’:* To generate a plot file for the execution of a contour or grey-scale plot for an image. It writes comments into the plot file.

*Task ‘LWPLA’:* To write the PostScript file.

PostScript is a programming language for describing how a page is to be displayed and this file is able to be inserted into the documents. It is similar to PDF. The file extension is either .ps or .eps.
Figure 9: Dirty image (all channels averaged) of G70, located in the centre of the plot.
Figure 10: CLEANED image (all channels averaged) G70, located in the centre of the plot
Figure 11: CLEANED image with self-calibration (all channels averaged) G70, located in the centre of the plot.
6. Data analysis

IDL is a programming language popularly used when one needs to process a large amount of data, for instance an image. Being vectorized, it can run vector operations as fast as a well-coded custom loop in fortran or C – languages from whom many constructs are inspired in IDL’s syntax. IDL runs more efficiently vectors, and so as any array programming language, one should smartly use the inbuilt vector operations to avoid slow processing time if elements are separately treated. It provides several mathematical analysis and graphical display techniques, and is usually a better and faster solution than the aforementioned languages when it comes to array-type data.

The H I data cube we retrieved from AIPS was too big to be managed by our software, which runs in IDL, so we cut it into a size of 10’ × 10’ containing G70 at the center of the cube.

We extracted H I absorption and emission spectra (Figure 13) from the radio continuum source G70. We selected a circular region around the source with the size of approximately 40” radius (about the size of the source itself) and averaged all pixels to create the absorption spectrum. This spectrum arises because the H I cloud absorbs the radio continuum at the wavelength of about 21 cm as mentioned in Chapter 3 of this report. For the emission spectrum, we selected a circular area around the source with a 64” radius. We then selected another area which is of the source size 40” radius), such as this one is embedded in the first area. Finally, we subtract this last area, which is like “cutting” the source out of the image, leaving a ring which is the one responsible for the emission spectrum. It expands about 25” away from the source. The brightness temperature, ΔT, axis is calculated by the Rayleigh-Jeans approximation and the velocity axis is obtained by Doppler’s shift formula and local standard at rest (LSR) velocity.

Both absorption and emission features are broadened due to H I moving with various velocities. The reasons for this are explained in Chapter 1 of this report.

These two spectra contain a wealth of information to investigate. To analyze the spectra, the assistance of a Gaussian fitting is needed. The Gaussian function is used to describe the normal distribution which is expressed in the form:

\[ A \exp\left(\frac{(-v - v_0)^2}{2\sigma^2}\right) \]

where \( A \) is a peak amplitude of Gaussian, \( v_0 \) is a mid-point velocity and \( \sigma \) is a standard deviation which usually control the width of the curve.
We first calculated the optical depth from equation (37) which is plotted in Figure 12. Then we fitted a Gaussian function to its components. The fitted parameters of the optical depth are shown in Table 1.

![Figure 12: Optical depth of H I cloud](image)

<table>
<thead>
<tr>
<th>$v [\text{km/s}]$</th>
<th>$A$</th>
<th>$\sigma [\text{km/s}]$</th>
</tr>
</thead>
<tbody>
<tr>
<td>-14.1</td>
<td>0.057</td>
<td>8.6</td>
</tr>
<tr>
<td>17.9</td>
<td>1.07</td>
<td>4.45</td>
</tr>
<tr>
<td>4.61</td>
<td>1.302</td>
<td>4.21</td>
</tr>
</tbody>
</table>

*Table 1: Fitted parameters of the optical depth*
To retrieve the spin temperature, $T_S$, and $HI$ column density, $N(HI)$, we again fitted Gaussians to both absorption and emission spectra which respectively yields $T_S$ by equation (38) and $N(HI)$ by equation (36) with the continuum level of 2320.17 K. The continuum temperature is calculated by Rayleigh-Jeans approximation. Derived and fitted parameters of the spectra are displayed in Table 2 and 3. In addition, we calculated a total $HI$ column density of $878 \times 10^{19} \text{cm}^{-2}$ by summing up all $HI$ column density in table 3.

Figure 13: Absorption (above) and emission (below) spectra toward G70. Negative velocity values refer to a movement toward us.

The $HI$ column density is computed from the emission spectrum, and not from the absorption one because there might be also emissions from the $HI$ cloud itself contributing inside the absorption line, whereas in the emission spectrum we subtracted already the continuum.
We searched for $H\ I$ at the location around G70 (Right ascension of 20 02 30 and declination of 33 30 30). The absorption spectrum shows a very broad absorption up to -40 km/s with the drop centered at about +3 km/s. This matches the measurements of recombination line (a bright line that produced by the combination process of an ion and an electron) at +3 km/s (Onello et al., 1995; Phillips et al., 1993), and might suggest that $H\ I$ originates from the recombination process. The ample well (wide range of velocities) indicates an expansion of $H\ I$ motion. The small absorption components at the velocities of about -52 km/s (Figure 17) and -70 km/s (Figure 18) are not real (not of our interest) but due to an irregularly-shaped cloud moving over the location of G70 in the sky.

The emission spectrum shows a wide emission at velocities around +5 km/s. As it can be seen, nearly all the emission lines are almost mirroring the absorption ones. This is easily understood as they belong to the same clouds (same velocity range).

Nevertheless, we have found no relationship between these emissions and the source. We also studied position-velocity diagrams created by collapsing one of the spatial axis of the cube (appendix A) but we could not see any interaction of G70 and ISM. An eventual interaction could have been detected if any significant shape of the $H\ I$ emission cloud would be revealed on either the channel map or position-velocity diagram, but this is not the case.

<table>
<thead>
<tr>
<th>$\nu_{abs}$ [km/s]</th>
<th>$A$ [K]</th>
<th>$\sigma$ [km/s]</th>
<th>$T_s$ [K]</th>
</tr>
</thead>
<tbody>
<tr>
<td>-12.15</td>
<td>-125.26</td>
<td>12.71</td>
<td>52.67</td>
</tr>
<tr>
<td>18.05</td>
<td>-1475.68</td>
<td>5.83</td>
<td>70.21</td>
</tr>
<tr>
<td>4.97</td>
<td>-1599.69</td>
<td>5.78</td>
<td>119.56</td>
</tr>
</tbody>
</table>

*Table 2: Fitted and derived parameters of the absorption spectrum*

<table>
<thead>
<tr>
<th>$\nu_{em}$ [km/s]</th>
<th>$A$ [K]</th>
<th>$\sigma$ [km/s]</th>
<th>$N(H\ I)$ [$10^{19}$ cm$^{-2}$]</th>
</tr>
</thead>
<tbody>
<tr>
<td>-4.16</td>
<td>57.58</td>
<td>2.57</td>
<td>64</td>
</tr>
<tr>
<td>19.96</td>
<td>64.06</td>
<td>2.68</td>
<td>74</td>
</tr>
<tr>
<td>12.28</td>
<td>105.95</td>
<td>2.26</td>
<td>103</td>
</tr>
<tr>
<td>5.78</td>
<td>111.45</td>
<td>13.33</td>
<td>637</td>
</tr>
</tbody>
</table>

*Table 3: Fitted and derived parameters of the emission spectrum*
Figure 14: H I emission around G70 at velocities of +16.18 to +26.47 km/s
Figure 15: HI emission around G70 at velocities of -2.39 to +5.87 km/s
Figure 16: H I emission around G70 at velocities of -25.05 to -18.86 km/s
Figure 17: H I emission around G70 at velocities of -55.99 to -45.67 km/s
Figure 18: \textit{H I} emission around G70 at velocities of -72.47 to -62.18 km/s
7. Conclusion and possible future work

The galactic source G70.7+1.2 is a hot topic among astronomers, as what we have been observing of it hardly matches our knowledge of celestial objects. In this framework, the scope of this study has been to reduce the data, derive key parameters (optical depth, spin temperature and column density) of the surrounding ISM, whose interaction with the source is used to improve our understanding of the object. The work is based on $H\ I$ data retrieved from the Westerbork Radio Telescope (WSRT).

We reduced the radio data using the dedicated package ‘Astronomical Image Processing System (AIPS)’. The reduction was done following a standard fashion in astronomy. After the UV fits files were read into AIPS, we applied a bandpass calibration using 3C286 and 3C147 because the observation was initially well calibrated in both phase and amplitude. The final steps consisted in imaging and self-calibration. Nevertheless, unwanted negative parts still remain post-reduction around the source, which is probably due to a difference between the theoretical and experimental dirty beam.

Subsequently, desired parameters have been solved and the spectra obtained. Unfortunately, the beam of the $H\ I$ data was too large to be useful for a more detailed analysis of the morphology and dynamics of the cold neutral ISM around G70.

Observations of the same object from $CO$ have been made with the James Clerk Maxwell Telescope (the largest single-dish telescope in the world which is designed to operate in sub-millimeter region), and multi-transition $CO$ data cubes have already been reduced. There is a relationship between the presence of $CO$ and molecular hydrogen. Consequently, a work to be envisaged as a next step to our study is the comparison of the atomic and molecular hydrogen (traced by $CO$), in order to identify interactions between them. Such a work had to be postponed due to stringent time limitations.


Appendix

A. PV-diagrams

*Figure 19: Position-velocity diagram of 15° rotated $H\ I$ data cube*
Figure 20: Position-velocity diagram of 30° rotated HI data cube

Figure 21: Position-velocity diagram of 45° rotated HI data cube
Figure 22: Position-velocity diagram of 60° rotated $\text{H I}$ data cube

Figure 23: Position-velocity diagram of 75° rotated $\text{H I}$ data cube
Figure 24: Position-velocity diagram of $90^\circ$ rotated $HI$ data cube

Figure 25: Position-velocity diagram of $105^\circ$ rotated $HI$ data cube
Figure 26: Position-velocity diagram of 120° rotated H I data cube

Figure 27: Position-velocity diagram of 135° rotated H I data cube
Figure 28: Position-velocity diagram of 150° rotated H I data cube

Figure 29: Position-velocity diagram of 165° rotated H I data cube
Figure 30: Position-velocity diagram of 180° rotated H I data cube