Probing the Galactic ISM in OH Absorption

By

Anita Petzler

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Examiner’s Copy
Except where acknowledged in the customary manner, the material presented in this thesis is, to the best of my knowledge, original and has not been submitted in whole or part for a degree in any university.

Anita Petzler
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Abstract

We examine the four $^2\Pi_{3/2}$, $J = 3/2$ ground state transitions of the hydroxyl radical (OH) along 15 sightlines through the Milky Way disk towards bright background continuum sources. We find that the OH gas along these sightlines is optically thin, consistent with the findings from the Southern Parkes Large Area Survey in Hydroxyl (SPLASH) pilot region [1]. We present a partially automated method for Gaussian decomposition of our spectra which identified 55 components across 28 OH-containing clouds. Three of these OH clouds had no associated $^{12}$CO emission (CO data from NANTEN [2]). We test two methods of finding excitation temperature: a modified ‘on-off’ method, and a method (outlined in [3]) where the excitation temperatures and column densities are solved for numerically. Neither method constrained $T_{\text{ex}}$ well. Column densities were calculated based on a range of possible excitation temperatures ($T_{\text{ex}}(1667) = 5-15$ K [4]), and conversions to $N$(H$_2$) were made. $N$(OH) values were found to range from $1 - 40 \times 10^{14}$ cm$^{-2}$, corresponding to $N$(H$_2$) values ranging from $1 - 40 \times 10^{21}$ cm$^{-2}$. We discuss our results in the context of CO-dark H$_2$, and outline future directions for our work.
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1 Introduction

1.1 Outline

A full census of the molecular interstellar medium (ISM) is essential to our understanding of how stars form. This thesis describes a study of the ISM within the disk of the Milky Way along 15 sightlines toward bright background continuum sources, via measurements of the hydroxyl radical (OH), a useful proxy for molecular hydrogen. Observations were made of the four $^2\Pi_{3/2}$, $J = 3/2$ ground state transitions of OH using the Australia Telescope Compact Array (ATCA). The details of these observations and subsequent data reduction are discussed in Chapter 2. The aims of this project were:

- Develop a partially automated, repeatable method to identify OH containing clouds within the Milky Way disk along the narrow ATCA synthesised beam. This aim was achieved through the Gaussian decomposition process described in Chapter 3, and facilitated the rest of the analysis in this work.

- Determine the peak optical depth of those clouds in the four lines of OH. This aim was partially motivated by analysis of the pilot region of the Southern Parkes Large Area Survey in Hydroxyl (SPLASH) [1]), which suggested that the OH lines were optically thin throughout the Galactic disk. Our higher resolution observations could determine whether this was due to beam-filling effects. Optical depth was determined via a simple radiative transfer treatment outlined in Section 1.3. The results are given in Chapter 4.

- Test different methods of determining the excitation temperature of the OH clouds. The first of these was a modified ‘on-off’ method where data from SPLASH was used to estimate the off-source spectrum. Second was a method whereby the excitation temperatures and column
densities of the four lines could potentially be solved for numerically. Both methods are described in Section 4.3. Unfortunately, the first method was limited in precision and the second was found to be inapplicable to this work. Instead, a ‘reasonable’ range of excitation temperatures (from [4] and references therein) was used for further analysis.

- **Determine the column densities of the OH clouds and use these to estimate the column density of H\textsubscript{2}.** The motivation is that OH column density can be converted to the total molecular hydrogen column density of the molecular component of the ISM, and that this may include diffuse material that is missed by CO. These results are presented in Chapter 4.

- **Compare our results to CO spectra (from the NANTEN Galactic Plane Survey [2]), to investigate whether we find evidence for CO-dark OH.** This was partially motivated by the lack of detections of OH without associated CO found in the SPLASH pilot region, and is discussed in Section 4.5.

The data used in this project was gathered as a compliment to SPLASH, but hold value in their own right as a probe of the Galactic ISM in the Inner Galaxy.

### 1.2 Cycle of Gas in Star Formation

Between the stars in our Milky Way Galaxy is a considerable amount of baryonic matter (dust and gas) which is collectively referred to as the interstellar medium (ISM) [5]. Hydrogen makes up the vast majority of the gas in the Galactic ISM, of which \( \approx 60\% \) by mass is atomic hydrogen (H\textsubscript{i}), \( \approx 23\% \) is ionised hydrogen (H\textsubscript{ii}) and \( \approx 17\% \) is molecular hydrogen (H\textsubscript{2}) [6]. The ISM also includes heavier elements, and the composition of the ISM at various Galactic radii is similar to the composition of the stars at those radii, as it is from the ISM that the stars form.

In broadest terms, the atomic gas of the ISM exists in two general temperature regimes: the cool neutral medium (CNM: \( T \approx 100 \text{ K, } n_{H} \approx 30 \text{ cm}^{-3} \))\(^1\) and the warm neutral medium (WNM: \( T \approx 5000 \text{ K, } n_{H} \approx 0.6 \text{ cm}^{-3} \)) [6]. Both phases are at the same pressure and are thermally stable, but gas at intermediate temperatures is thermally unstable. Therefore, if perturbed out of equilibrium, collisionally excited particles within the WNM can rapidly de-excite radiatively, cooling to join the CNM. The exact mechanisms of these processes depend on a number of complex factors (e.g. composition of the gas, magnetic fields, turbulence, etc.), the overall effect being that the atomic gas of the ISM will condense and cool, eventually allowing molecular hydrogen to form and persist [1, 7–9].

\[^1\text{n}_{H} = \text{Number density: the number of hydrogen atoms per cubic centimetre.}\]
Focussing on this molecular gas, it will initially form in the shielded interior of condensing gas clumps, as the extinction in these regions will be sufficient to protect the H$_2$ from the dissociating radiation that exists throughout the ISM. A significant portion of the molecular gas at this stage could still be quite diffuse, particularly if the cloud that contains it has a relatively low mass [10]. As will be discussed in Section 1.4, this more diffuse molecular gas is not reliably traced by carbon monoxide emission – the most commonly used tool for detecting molecular gas – whereas the more dense gas is. OH is expected to coincide with H$_2$ through all density regimes [11, 12]. As such, measurements of OH provide valuable, additional information to that provided by observations of CO.

1.3 Radiative Transfer

In astronomy our ability to gather information about the objects we study is limited to what we can infer from the radiation we receive. In this section we discuss the physical principles that allow us to infer the local conditions of a gas cloud from spectral line observations.

Our measurements of intensity take advantage of the fact that in the radio regime, Planck’s Law relating the intensity and temperature of a blackbody (Equation 1.1) is approximated well by the Rayleigh-Jeans Law (Equation 1.2), which is linear:

\[
B_\nu(T) = \left(\frac{2h\nu^3}{c^2}\right) \frac{1}{e^{h\nu/k_BT} - 1},
\]

\[
B_\nu(T) = \frac{2\nu^2k_BT}{c^2}.
\]

If we solve Equation 1.2 for temperature we can define the quantity ‘brightness temperature’,

\[
T_b = \frac{I_\nu c^2}{2\nu^2k_B},
\]

which is the temperature of a blackbody that emits radiation of intensity $I_\nu$ at frequency $\nu$.

We also express the relative populations of the energy levels of a spectral line transition by the excitation temperature, $T_{ex}$ defined by the Boltzmann factor (Equation 1.4):

\[
\frac{n_u}{n_l} = \frac{g_u}{g_l} e^{-\frac{h\nu_0}{k_BT_{ex}}},
\]

The response of radio telescope antennas to incoming radiation can be calibrated to relate directly to these observable ‘temperatures’.
1.3.1 Spectral line observations

When an atomic or molecular system transitions from a higher energy state to a lower energy state, a photon is emitted. That photon will have energy equal to the difference in energy of the two states. If the particles in a gas undergo many of these same transitions, these photons will be collectively seen as a spectral emission line at the frequency corresponding to the energy of those photons. This process is dependent upon the relative population of the higher and lower energy states. In the ISM the initial excitation is generally produced collisionally (though it can be produced radiatively), often between the species in question and hydrogen. Thus when a spectral emission line is observed we can make the broad statements that the element or molecule associated with the transition is present along the line of sight, and that local conditions (i.e. temperature, density) are sufficient to significantly populate the higher level of the transition. Similarly, the system may transition from a lower energy state to a higher energy state by absorbing a photon, which can result in an absorption line. For an absorption line to be produced, there must generally exist a background radiation source emitting sufficiently bright radiation at the appropriate frequency.

If there is relative motion between the observer and the emitting/absorbing gas, the frequency at which the spectral lines are seen will be Doppler shifted. This relative motion can be bulk motion of the gas which leads to a shifted line, or motion within the gas (e.g. thermal motion, turbulence, velocity gradients, etc.) which leads to a broadened line. This shifting and broadening is in frequency, but provided that the rest frequency is known, these frequency shifts can be translated back to line-of-sight velocity, resulting in velocity spectra. Both thermal and turbulent broadening result in Gaussian-shaped line profiles in velocity\(^2\), which can be individually identified and fit in the spectrum\(^3\). A Gaussian line profile is defined by:

\[
\tau_\nu(\nu) = \tau_{\text{peak}} e^{-4 \ln 2 (c - \nu_0)^2 / \Delta \nu^2}.
\]

(1.5)

Note that in this case we have expressed the line profile in terms of optical depth \(\tau_\nu\) (defined in Section 1.3). The Gaussian profile is described by its mean velocity \(\nu_0\), its peak optical depth \(\tau_{\text{peak}}\) and its full width at half-maximum \(\Delta \nu\).

The degree to which each line is shifted or broadened depends directly on the magnitude and dispersion of radial velocity of the gas relative to the observer. Thus when a specific spectral line is

\(^2\)Thermal broadening is easily shown to be Gaussian via Maxwell’s velocity distribution equation. Turbulent broadening is much more complex, but the turbulent motions are generally accepted to be approximately Maxwellian, and hence Gaussian.

\(^3\)When multiple blended lines exist the process of Gaussian decomposition is highly non-trivial, as discussed further in Section 3.
observed, a range of frequencies both higher and lower than the target transition line are included in order to ‘catch’ gas in an expected velocity range. When spectral line emission and absorption through the plane of our Galaxy are observed, as is done in this project, the bulk velocities arise primarily from the differential rotation of the Galactic disk. Thus the velocity of an emitting or absorbing cloud can be used to determine its likely position in the Galaxy.

1.3.2 Spectral line absorption and emission

As radiation of intensity $I_\nu$ passes a distance $ds$ along the line of sight, it will be modified by that gas via a process of radiative transfer:

$$\frac{dI_\nu}{ds} = -\kappa_\nu I_\nu + j_\nu,$$

where $\kappa_\nu$ and $j_\nu$ are the spectral absorption and emission coefficients, respectively. In a uniform slab of gas\(^4\) the absorption and emission coefficients are defined by:

$$\kappa_\nu = \frac{c^2}{8\pi\nu^2_0} \frac{g_u}{g_l} n_l A_{ul} \left(1 - e^{-\frac{\hbar \nu}{k_B T_{ex}}} \right) \phi(\nu),$$

$$4\pi j_\nu = n_u h \nu A_{ul} \phi(\nu),$$

where $n_u$ and $n_l$ are the number densities of the upper and lower levels of the transition, $\phi_\nu$ is the normalised line profile $\int_{-\infty}^{\infty} \phi(\nu) d\nu = 1$ and $A_{ul}$ is the Einstein A coefficient.

Integrating Equation 1.7 over line-of-sight distance will then change the absorption coefficient $\kappa_\nu$ into the dimensionless quantity optical depth $\tau_\nu$ (as $\kappa_\nu ds = d\tau_\nu$), and the number densities $n$ into column densities $N$, defined as the number of particles per unit area along the line of sight. This then gives us an equation for optical depth:

$$\tau_\nu = \frac{c^2}{8\pi\nu^2_0} \frac{g_u}{g_l} N_l A_{ul} \left(1 - e^{-\frac{\hbar \nu}{k_B T_{ex}}} \right) \phi(\nu).$$

The value of optical depth $\tau_\nu$ will depend on the composition, density, excitation state and physical dimensions of the gas, as well as the frequency at which it is observed. It is common for gas clouds in the ISM to have an optical depth much less than one when viewed in the radio spectrum. These clouds are generally described as ‘optically thin’, whereas clouds with optical depth much greater than one are described as ‘optically thick’. Optical depth can also be negative, implying an amplification of the incident radiation.

\(^4\)Molecular clouds are known to have clumpy or filamentary structure on all observable spatial scales [13, 14], so this assumption will lead to an ‘averaging’ of this structure.
1.3 Radiative Transfer

In the radio regime $\hbar \nu_0 \ll k_B T_{\text{ex}}$, allowing us to retain only the first term in the series expansion of the exponential. Integrating Equation 1.9 over frequency and rearranging then results in an equation for column density of the lower level of the transition, $N_l$:

$$N_l = \frac{8\pi \nu_0}{c^2} \frac{g_l k_B T_{\text{ex}}}{g_u h A_{ul}} \int \tau_\nu d\nu.$$  \hspace{1cm} (1.10)

This expression can be converted from units of frequency to velocity via the relation $d\nu = d(3\nu_0 c)$. $N_l$ can be converted to the total column density of the species $N$ via the following relation:

$$\frac{N_l}{N} = \frac{g_l e^{-E_l/k_B T_{\text{ex}}}}{\sum g_i e^{-E_i/k_B T_{\text{ex}}}},$$

which assumes that the populations of all energy levels may be described by a single excitation temperature, $T_{\text{ex}}$. In the case of the ground-state OH lines considered in this work, we may assume that the factor $\hbar \nu/k_B T_{\text{ex}} \ll 1$, and:

$$\frac{N_l}{N} \approx \frac{g_l (1 - \hbar \nu_{ul}/k_B T_{\text{ex}})}{\sum g_i (1 - \hbar \nu_{ij}/k_B T_{\text{ex}})} \approx \frac{g_l}{\sum g_i},$$

which results in the following expression for the OH column density, once the degeneracies of the ground state levels are substituted (see Section 1.5):

$$N(\text{OH}) = \frac{8\pi \nu_0^2}{c^3} \frac{16 k_B T_{\text{ex}}}{g_u h A_{ul}} \int \tau_\nu d\nu.$$  \hspace{1cm} (1.13)

We note that this expression also assumes that only the ground state level of OH is populated, which is a reasonable assumption for most ISM conditions.

Returning now to the equation of radiative transfer (1.6), if we define a source function $S_\nu \equiv \frac{j_\nu}{\kappa_\nu}$, it can be shown from the definitions of $j_\nu$ and $\kappa_\nu$ that:

$$S_\nu = \left(\frac{2\hbar \nu_0^3}{c^2}\right) e^{\hbar \nu_0/k_B T_{\text{ex}}} - 1 - 1$$

$$= B_\nu(T_{\text{ex}}),$$

By substituting Equation 1.14 into Equation 1.6 and solving the differential equation, we obtain the solution to the equation of radiative transfer:

$$I_\nu(\tau_\nu) = I_\nu(0) e^{-\tau_\nu} + B_\nu(T_{\text{ex}})(1 - e^{-\tau_\nu}),$$

where $I_\nu(\tau_\nu)$ is the observed intensity and $I_\nu(0)$ is the background continuum intensity. In the radio regime the Planck Function $B_\nu$ is approximated by the Rayleigh-Jeans Law, meaning that we may relate
$I_\nu$ linearly to the brightness temperature, and $B_\nu(T_{\text{ex}})$ linearly to $T_{\text{ex}}$ according to the same expression. The intensity terms can therefore be written in terms of their associated temperatures, resulting in:

$$T_b^* = T_c e^{-\tau_\nu} + T_{\text{ex}} (1 - e^{-\tau_\nu}).$$

The background continuum brightness temperature $T_c$ can be subtracted from this observed brightness temperature $T_b^*$ to give the continuum subtracted brightness temperature $T_b$:

$$T_b = (T_{\text{ex}} - T_c) (1 - e^{-\tau_\nu})$$

(1.16)

The expected continuum-subtracted brightness temperatures observed for different combinations of excitation temperature and optical depth are outlined in Figure 1.1.

### 1.4 Tracers of Molecular Hydrogen

Due to the symmetry and low mass of the H$_2$ molecule even its lowest lying transitions are not typically excited at the temperatures seen in much of the molecular ISM, particularly in the cold interiors of clouds. Molecular hydrogen in the ISM is therefore often studied via the spectral lines of other species that co-exist with H$_2$; some examples relevant to this discussion are carbon monoxide (CO), hydroxyl (OH) and ionised carbon (C$^+$). This work examined the ISM along sightlines through the Galactic disk which has complicated structure along the line of sight. Spectral line tracers such as CO, OH and C$^+$ are therefore extremely useful for their ability to disentangle spatially coincident features by their differing velocities. This is in contrast to tracers such as broad-band dust emission, which despite some key advantages, does not give this crucial velocity information.

The emission of the CO $J = 1 \rightarrow 0$ transition at 115.3 GHz has been used extensively as a proxy for H$_2$ since the 1980s [15, 16], with an empirical conversion factor often used to convert CO integrated intensities directly to molecular hydrogen column densities. Though the conversion from CO to $N$(H$_2$) has high uncertainties [17], its relative ease of observation and analysis allow CO to remain a useful and popular tool. Nevertheless, even during the earliest days of its use, CO was expected to not trace a significant amount of molecular gas in diffuse regions [18, 19] where the extinction of the cloud is sufficient to shield H$_2$ but not CO [20, 21]. However, the regions of most interest were generally sufficiently massive and dense that this limitation was not significant [10].

In recent decades our exploration has focussed more and more on these previously neglected diffuse regions, leading to increased interest in species other than CO that are expected to trace diffuse molecular gas. Hydroxyl (OH) is expected to form in the same environments as molecular hydrogen,
1.4 Tracers of Molecular Hydrogen

(a) Net absorption:
\[ \tau_\nu > 0, \quad 0 < T_{\text{ex}} < T_c \]

(b) Net emission (not masing):
\[ \tau_\nu > 0, \quad T_{\text{ex}} > T_c \]

(c) Net emission (masing):
\[ \tau_\nu < 0, \quad T_{\text{ex}} < 0 \]

Figure 1.1: As light from a bright continuum emitting source travels towards the observer, it will pass through the gas of the interstellar medium. This figure considers the effect on the observed brightness temperature of two parameters: optical depth \( \tau_\nu \) and excitation temperature \( T_{\text{ex}} \). (a) If \( \tau_\nu \) is positive and \( 0 < T_{\text{ex}} < T_c \) then there will be net absorption, and the observed brightness temperature will be less than the continuum brightness temperature. (b) If \( \tau_\nu \) is positive and \( T_{\text{ex}} > T_c \) there will be net emission from the cloud as the excitations into the upper level of the transition caused by light from the background continuum source are outnumbered by the total de-excitations into the lower state. (c) If both \( \tau_\nu \) and \( T_{\text{ex}} \) are negative then the light from the background continuum source will be amplified by the cloud, resulting in the observation of an astrophysical maser.
1.5 OH Ground State and its Lines

and has been demonstrated to trace diffuse, so-called ‘CO-dark’ H$_2$ [22–26]. Observations of the four $^2\Pi_{3/2}$, $J = 3/2$ ground state transitions of the hydroxyl molecule can be used to determine the column density of OH, which is then used as a tracer of H$_2$. OH column density is linearly correlated with 100 $\mu$m dust emission [27] implying that it traces the same gas as the dust, but with the added velocity information of a spectral tracer.

This highlights the ability of OH to provide information supplementary to that given by measurements of CO, and motivates much of the present work. Other spectral line tracers of CO-dark H$_2$ include ionised and neutral carbon (C$^+$, C respectively). An example is the Galactic Observations of Terahertz C$^+$ survey (GOT C$^+$, [28, 29]), which used C$^+$ emission along a sparse grid of sightlines to trace molecular gas in the Galactic plane [30, 31]. With its focus on OH, this work is complimentary to studies of C$^+$, such as GOT C$^+$.

### 1.5 OH Ground State and its Lines

The OH molecule consists of one oxygen atom and one hydrogen atom bound by a single covalent bond. This leaves one unpaired valence electron from the oxygen atom. The $^2\Pi_{3/2}$, $J = 3/2$ ground state of hydroxyl is split into four levels, shown in Figure 1.2. The four levels in the ground state arise from interactions between the rotation of the molecule and the orbit of the unpaired electron ($\Lambda$-doubling) and the hyperfine interaction between the spins of the electron and the nuclei [32].

The four ground state levels have very small energy gaps, and transitions between the levels can absorb or emit radio-frequency photons at rest frequencies of 1612.231, 1665.402, 1667.359 and 1720.530 MHz. The 1665 and 1667 MHz transitions are commonly referred to as the main lines, while the 1612 and 1720 MHz transitions are referred to as the satellite lines [33]. In order of increasing rest frequency, the ratio between the quantum mechanical transition strengths is 1 : 5 : 9 : 1. If the excitation temperature of all four transitions is identical, this results in observed brightness temperature ratios of 1 : 5 : 9 : 1 in the optically thin limit and 1 : 1 : 1 : 1 in the optically thick limit [34]. However, from the earliest days of observations [35–39] it was clear that anomalous excitation of the ground state transitions of OH was common: the transitions were not in local thermodynamic equilibrium (LTE) and hence had different excitation temperatures. This means that we cannot make naive assumptions about $T_{ex}$. On the other hand, the four transitions share the same four energy levels, meaning that their excitations are not fully independent. This section will discuss the mechanisms behind the anomalous excitation of the four OH ground state lines, outline the relationships expected to exist between them, and present examples of typical observed line profiles.
1.5 OH Ground State and its Lines

Figure 1.2: From Figure 1 in Dawson et al. (2014) [1]. Energy level diagram showing that the $^2\Pi_{3/2}, J = 3/2$ ground state of hydroxyl is split into four levels due to $\Lambda$-doubling and hyperfine splitting. The transitions shown represent the emissions at 1612.231, 1665.402, 1667.359 and 1720.530 MHz. $F$ is the total angular momentum quantum number.

1.5.1 OH excitation – rotational ladder

The anomalous excitation seen in the ground state lines can be understood by considering population of the ground state via downward transitions from higher rotational levels, as shown in Figure 1.3. Excitation into rotational states above the $^2\Pi_{3/2}, J = 3/2$ ground state will cascade back to the ground state via the $^2\Pi_{3/2}, J = 5/2$ intra-ladder rotational state or the $^2\Pi_{1/2}, J = 1/2$ cross-ladder rotational state [33]. Here, it is only necessary to consider excited rotational states, as electronic and vibrational excitations require more energy than is commonly available in the ISM [33]. If the transitions between the $^2\Pi_{3/2}, J = 3/2$ ground state and the first two excited rotational states are optically thick\(^5\), the number of possible transitions into a given ground state level is the only factor that determines their relative populations [33].

As can be seen in Figure 1.3 cascades down to the ground state via the $^2\Pi_{3/2}, J = 5/2$ state will tend to over-populate the $F = 2$ levels of the ground state. If this cascade path dominates the population of the ground state levels then it would cause a population inversion of the 1720 MHz line and hence give that line a negative excitation temperature. This would also lead to a sub-thermal excitation temperature in the 1612 MHz line due to the over-population of its lower level. Likewise, cascades down to the ground state via the $^2\Pi_{1/2}, J = 1/2$ state will tend to over-populate the $F = 1$ levels of the ground state. This will lead to a negative excitation temperature in the 1612 MHz line.

\(^5\)This is not always the case: deviations from this condition will be discussed later in this section.
Figure 1.3: Schematic of the three lowest rotational states of OH, as well as the lambda and hyperfine splitting of those states. Excitations above the $^2\Pi_{3/2}$, $J = 3/2$ ground state will cascade back down to it via the $^2\Pi_{3/2}$, $J = 5/2$ intra-ladder level, or the $^2\Pi_{1/2}$, $J = 1/2$ cross-ladder level. The statistical weight of each level is given by the overall angular momentum quantum number $F$ via: $2F + 1$. Allowable transitions are those where parity is changed and $\Delta F = 1, 0$. These are shown in green and red on the schematic.

and a sub-thermal excitation temperature in the 1720 MHz line. In both of these scenarios the main lines (1665 MHz: $F = 1 \rightarrow 1$; 1667 MHz: $F = 2 \rightarrow 2$) would not be affected by this anomalous excitation and would be expected to have excitation temperatures approximately equal to one another. These would typically lie between the radiation temperature in diffuse regions ($\approx 3$ K) and the kinetic temperature in more dense, thermalised regions ($\approx 15$ K).

This relationship between the excitation temperatures of the ground state lines is quantified by the excitation temperature sum rule, which is obtained directly from the definition of excitation temperature (1.4):

$$\frac{\nu_{1665}}{T_{\text{ex}(1665)}} + \frac{\nu_{1667}}{T_{\text{ex}(1667)}} = \frac{\nu_{1612}}{T_{\text{ex}(1612)}} + \frac{\nu_{1720}}{T_{\text{ex}(1720)}}. \quad (1.17)$$

Equation 1.17 can therefore be used to constrain possible excitation temperature values.

The degeneracies, $g$, of the hyperfine sub-levels are given by $2F + 1$. As described by [33], the $2F + 1 = 3$ sub-levels of the $F = 1$ level are more easily overpopulated than the $2F + 1 = 5$ sub-levels of the $F = 2$ level, so the population inversion of the 1612 MHz line naturally dominates that of
1.5 OH Ground State and its Lines

the 1720 MHz line. In order for the inversion of the 1720 MHz line to occur, cascades through the $^2\Pi_{1/2}, J = 1/2$ state must be suppressed [33]. This can be achieved in two ways:

1. If the $^2\Pi_{1/2}, J = 1/2$ transitions are optically thin and the $^2\Pi_{3/2}, J = 5/2$ transitions are optically thick.

2. If excitations to any states higher than $^2\Pi_{3/2}, J = 5/2$ are suppressed.

The first condition can be met in environments where the OH column density is relatively low, since the cross-ladder transition strengths are an order of magnitude lower than the intra-ladder transition strengths [33]. This condition was examined closely by van Langevelde et al. [39]. They found that the $^2\Pi_{3/2}, J = 5/2$ transition becomes optically thick at $N_{\text{OH}}/\Delta V \approx 1 \times 10^{14} \text{ cm}^{-2} \text{ km s}^{-1}$, where the $^2\Pi_{1/2}, J = 1/2$ transition becomes optically thick at $N_{\text{OH}}/\Delta V \approx 9 \times 10^{14} \text{ cm}^{-2} \text{ km s}^{-1}$. Thus in the $1 \times 10^{14} < N_{\text{OH}}/\Delta V < 9 \times 10^{14}$ regime we would expect to see an inversion in the 1720 MHz line, and an inversion in the 1612 MHz line when $N_{\text{OH}}/\Delta V > 9 \times 10^{14}$. The second condition is met in cool environments dominated by collisions, but is almost impossible in environments dominated by radiation [33].

It is also common to see some degree of anomalous excitation in the main lines, often evidenced in observations as a ‘non-LTE’ ratio of main-line brightness temperatures. Here LTE stands for local thermodynamic equilibrium, and is usually defined as a situation in which collisions dominate the populations of the levels of interest, such that they are described by a single $T_{\text{ex}}$ that is equal to the gas kinetic temperature. In the case of LTE, the brightness temperatures of the main lines are expected to have the ratio $T_b(1667)/T_b(1665) = 1.8$ in the optically thin regime, and 1.0 in the optically thick regime. Many observers (e.g. [40], [1] and references therein) will use these ratios as a first indicator of whether the gas might be in LTE. If this ratio holds, it has historically been common to assume that both main lines share a single $T_{\text{ex}}$ and state that they are ‘in LTE’ even if the satellite lines are anomalously excited [1, 41]. This characterisation can be useful but is also misleading. The main lines of OH can indeed have the same excitation temperature, but the two separate pairs of levels that produce them will usually contain one overpopulated pair and one underpopulated pair, as described above. This becomes problematic when main line brightness temperature ratios are assumed to imply true LTE, with this assumption then used to estimate optical depths and column densities (as done by e.g. [40, 42]. As pointed out by [43] and [1], this leads to significant errors in the derived quantities. In fact, the main-line brightness temperature ratio can readily fall in the permitted ‘LTE’ range of 1.0 – 1.8 even if the main-line excitation temperatures are not equal. Therefore, though the literature
makes extensive use of the term ‘LTE’ in reference to the main lines of OH, we will instead be more explicit and refer to their excitation temperatures.

1.5.2 Typical profiles

The brightness temperature spectrum of an OH cloud will be determined by the solution to the radiative transfer equation seen in Equation 1.16, and therefore depends on the excitation temperature $T_{\text{ex}}$, the brightness temperature of the background continuum $T_c$ and the optical depth $\tau_\nu$. In general, emission in any of the four OH ground state lines will occur when $T_{\text{ex}}$ (and hence $\tau_\nu$) is negative if $T_c > T_{\text{ex}}$, and when $T_{\text{ex}}$ (and hence $\tau_\nu$) is positive if $T_c < T_{\text{ex}}$. Figure 1.4 shows velocity profiles of the four ground state transitions toward a bright [39] and a diffuse [1] background continuum.

1.6 Radio Interferometry

The discussion of radio interferometry in this section follows that of Avison and George (2013 [44]). As will be discussed in more detail in Chapter 2, one of the aims of our observations was to resolve the ambiguity of the low optical depths inferred from analysis of the SPLASH pilot region [1]. If these low optical depths were the result of averaging over the Parkes radio telescope’s $\sim 12$ arcmin beam, then a higher resolution investigation of this region should yield higher optical depths along some sightlines. The resolution ($\theta$/arcmin) of a filled-aperture telescope like Parkes varies as $\theta = 1.22 \lambda/D$, where $\lambda$ is the observing wavelength and $D$ is the diameter of the dish. In order to increase our resolution by a factor of 24 (as presented in this work) a single dish $> 1.5$ km in diameter would be required. This is more than twice as large as the world’s largest single-dish radio telescope⁶. Instead, radio astronomers can use interferometry to achieve high resolution using an array of smaller radio telescopes, while sacrificing sensitivity. The maximum theoretical resolution of an interferometer is given by $\theta \approx \lambda/B_{\text{max}}$ where $B_{\text{max}}$ is the maximum separation (baseline) of the antennas. A schematic of a simple two dish interferometer is shown in Figure 1.5 and the discussion that follows focuses on this simple system.

As wavefronts from the observed source arrive at each antenna, a voltage is induced in the receiver that is related directly to the brightness of the source, the area of the dish and the solid angle of the sky ‘seen’ by the telescope. The separation of the two antennas introduces a geometric time delay $\tau_{\text{geo}}$ at

⁶RATAN-600 is a 600 m diameter partially-filled single aperture radio telescope in Zelenchukskaya, Russia. The largest fully-filled aperture telescope is the Five-hundred-meter Aperture Spherical radio Telescope (FAST) in Guizhou, China.
Figure 1.4: The panel at top is from Figure 1 in van Langevelde et al. (1995 [39]) and shows velocity profiles of the four $^2\Pi_{3/2}, J = 3/2$ ground state transitions toward the nucleus of Centaurus A: a bright background continuum source. Note that these profiles show flux vs heliocentric velocity and have not been continuum subtracted. The conjugate nature of the anomalous excitation of the satellite lines is clearly demonstrated by the lowest trace which shows that the sum of the satellite lines tends to a constant value. The panel at bottom is from Figure 3 in Dawson et al. (2014 [1]) and shows velocity profiles of the four $^2\Pi_{3/2}, J = 3/2$ ground state transitions toward G342.40+0.35 where the background continuum brightness temperature was approximately 17 K. Based on the discussion in Section 1.5.1, the emission seen in the top panel is indicative of a negative excitation temperature. On the other hand, the main line emission in the bottom panel (at $v = -140$ km s$^{-1}$) may indicate that their excitation temperatures are greater than the background continuum brightness temperature.
1.6 Radio Interferometry

Figure 1.5: Two antennas \( x \) and \( y \) with a baseline separation of \( \mathbf{b}_\\lambda \) observe a source located along the unit vector \( \mathbf{s} \) at an elevation of \( \rho \). Wavefronts from the source (shown in blue) are received at antenna \( x \) first, then at antenna \( y \) after a geometric time delay of \( \tau_{\text{geo}} \). This schematic is taken from figure 1 in Avison and George (2013) [44].

which those wavefronts arrive that is given by:

\[
\tau_{\text{geo}} = \frac{\mathbf{b} \cdot \mathbf{s}}{c} = \frac{bs \cos \rho}{c},
\]  

(1.18)

where \( \mathbf{b} \) is the baseline vector, \( \mathbf{s} \) is the unit vector in the direction of the object to be observed, \( \rho \) is the elevation of the object and \( c \) is the speed of light.

The signals received by the two antennas are then cross correlated by a computer, resulting in an output \( R_{xy} \) equivalent to the multiplication of the Fourier transform of the signal from one antenna \( X(\nu) \) with the complex conjugate of the transform of the signal from the other antenna \( Y^*(\nu) \). This output is related to the source intensity, \( I(\mathbf{s}) \), by:

\[
R_{xy}(\tau_{\text{geo}}) = X(\nu) Y^*(\nu) = \Delta \nu \int A(\mathbf{s}) I(\mathbf{s}) \cos(2\pi \mathbf{b}_\\lambda \cdot \mathbf{s}) d\Omega,
\]  

(1.19)

where \( A(\mathbf{s}) \) is the total area of the dishes perpendicular to the line of sight \( \mathbf{s} \), and the baseline length is measured in units of the observed wavelength \( \lambda \). The range of frequencies observed, or bandwidth is given by \( \Delta \nu \).

If we normalise this based on the phase tracking centre, \( \mathbf{s}_0 \), where all other sky positions are described by their relative offset \( \mathbf{\sigma} \), such that \( \mathbf{s} = \mathbf{s}_0 + \mathbf{\sigma} \), we can define a ‘complex visibility’ \( V \):

\[
V = |V| e^{-i\phi} = \int A(\mathbf{\sigma}) I(\mathbf{\sigma}) e^{-i2\pi \mathbf{b}_\\lambda \cdot \mathbf{\sigma}} d\Omega,
\]  

(1.20)
where $A(\sigma)$ is the projected area of the dishes, with their responses normalised by dividing by a factor of $A_0 = A(s_0)$. Equations 1.19 and 1.20 can be combined to form:

$$R_{x,y} = A_0 |V| \Delta v \cos(2\pi b \cdot s_0 - \phi_0).$$  \hspace{1cm} (1.21)

Thus the complex visibility $V$ can be determined from the Fourier transform of the output from the correlator from Equation 1.21, which can in turn give a component of the sky brightness distribution $I(\sigma)$ from Equation 1.20. Hence we are able to reconstruct the sky brightness distribution from the complex visibility.

A considerable simplification occurs if the coordinate system is re-arranged so that measurements of $V$ are taken on a plane. The unit vector $s$ is defined by its projection onto a Cartesian coordinate system with axes $(u, v, w)$, co-moving with the source of interest. For small angles, only the $u$ and $v$ coordinates need be considered and hence the $uv$-plane is the two-dimensional projection of the antenna baselines onto the sky. In that coordinate frame, Equation 1.20 is transformed to:

$$V = \int A(l, m) I(l, m) e^{-i 2\pi (ul + vm)} \frac{dl \, dm}{\sqrt{1 - l^2 - m^2}}.$$  \hspace{1cm} (1.22)

where $l$ and $m$ are projections of the vector $s$ in the east-west $(u)$ and north-south $(v)$ directions, respectively. If both $l$ and $m$ subtend small angles, the small angle approximation can be used, allowing $l$ and $m$ to be replaced by their angular offsets from the phase centre, $i$ and $j$. Therefore the complex visibility $V(l, m)$ is the Fourier transform of the sky brightness distribution $I(i, j)$.

A fundamental property of the Fourier transform means that short baselines correspond to large scales on the sky, while long baselines correspond to small angular scales. In order to recover $I(i, j)$ at all spatial scales, in theory we must sample $V(l, m)$ at all $uv$-spacings. In practice this is not always necessary, but we must have sufficient coverage at those spacings corresponding to our spatial scales of interest. The resolution of the telescope, and hence the size of the synthesized beam, is determined by the largest $uv$-spacings while the largest spatial scales to which the observations are sensitive are determined by the shortest $uv$-spacings. Incomplete sampling of the $uv$-plane between these limits will result in a loss of sensitivity at the corresponding spatial scales which can introduce artifacts into the final images.

### 1.6.1 Caveats of interferometer observations

A caveat of using an interferometer (e.g. the ATCA) rather than a single dish (e.g. Parkes), is that due to the spatial filtering caused by an incomplete sampling of the $uv$-plane, the interferometer will not be sensitive to flux on all spatial scales [45]. In Figure 1.67, we assume that the OH gas between a compact
Figure 1.6: Illustration of the spatial filtering effect of interferometer observations. The graphic at top illustrates the relative size on the sky of the interferometer (ATCA) synthesised beam, an intervening gas cloud and a background continuum emitting source. Image (a) represents the observer’s view of the background continuum source and the intervening gas cloud. (Note that we have set the emission from the continuum three times brighter than that from the cloud for clarity in the illustration, but the continuum sources examined in this work are in fact orders of magnitude brighter.) Image (b) shows the image of this cloud and continuum source produced from the ATCA using the same array configuration as the observations in this project. Note that while the background continuum source is well represented in the observed image, most of the flux from the intervening cloud is lost. As a comparison, image (c) shows the same background continuum source without the intervening gas cloud. The observed image (d) of this source is similar to the image with the intervening cloud (b). Images (a-d) were produced using the Friendly Virtual Radio Interferometer software (https://crpurcell.github.io/friendlyVRI/).
continuum source and the observer is much larger on the sky than the largest spatial scales to which our observations are sensitive, which is determined by the shortest $uv$-spacings. This uniformity means that the interferometer will be unable to detect flux from the diffuse OH gas. However, as previously noted, if the continuum source is compact and/or of a comparable angular diameter to the synthesised beam (determined by the largest $uv$-spacings) then flux from the continuum (and hence absorption against the continuum) will be well-recovered. In order to see the effect of this, we take another look at Equation 1.16:

$$T_b = (T_{ex} - T_c)(1 - e^{-\tau})$$

If we rearrange this equation to form Equation 1.23, we can see that the continuum subtracted brightness temperature $T_b$ depends on two factors: emission from the large OH gas cloud, represented by the $T_{ex}$ term, and absorption of the continuum flux, represented by the $T_c$ term:

$$T_b = T_{ex}(1 - e^{-\tau}) - T_c(1 - e^{-\tau}).$$

However, if the OH cloud is smooth and extended compared to the spatial scales to which the interferometer is sensitive, the flux detected from the $T_{ex}$ term will be negligible, and it can be omitted from the equation of radiative transfer [45, 46]. We are then left with Equation 1.24 which can be readily solved for optical depth from our observations of $T_b$ and $T_c$:

$$T_b = -T_c(1 - e^{-\tau}).$$

Furthermore, we note here that even if our observations included zero $uv$-spacings (i.e. sensitivity to the largest spatial scales), provided that the brightness of our continuum sources is much greater than the expected excitation temperatures, the $T_c$ term in Equation 1.23 dominates, and Equation 1.24 still provides a very good estimation of $T_b$. Very bright continuum sources were chosen for this project partially for this reason, mitigating against the likely possibility that the OH clouds had clumpy or filamentary structure on the scales to which our observations were sensitive.

### 1.7 Bayesian Statistics

Bayesian inference is an approach to statistics that incorporates knowledge about prior information and about how ‘sensible’ the data are. In the context of this project, Bayesian statistics will be applied

\footnote{Note that Figure 1.6 (a) shows a background continuum source that is less bright when compared to the emission of the intervening cloud than those examined in this work.}
1.7 Bayesian Statistics

Bayesian Statistics to the process of Gaussian decomposition described in Chapter 3. Specifically, a Bayesian approach will be used to evaluate competing Gaussian models of our spectral data.

For a model hypothesis \( M \) and data \( d \), the posterior probability for a set of model parameters \( \theta \) is given by Bayes’s Theorem:

\[
P(\theta | d, M) = \frac{P(d | \theta, M) P(\theta | M)}{P(d | M)}.
\] (1.25)

Here \( P(\theta | d, M) \) is the probability that a model \( M \) with parameters \( \theta \) fits a set of data \( d \). This ‘posterior probability’ is informed by the ‘likelihood’ \( P(d | \theta, M) \) (i.e., how well the data fits the model), and the ‘prior’ information about the probability of the model parameters \( P(\theta | M) \), as well as the probability of the data given the model \( P(d | M) \). If no prior information is available, then this method is equivalent to maximum-likelihood analysis. If the uncertainties on the data are well described by Gaussian statistical noise, then the likelihood is directly related to the \( \chi^2 \) statistic via \( P(d | \theta, M) = -\chi^2/2 \).

Once several models have been proposed and optimised to fit a dataset, then the preferred model can be chosen by calculating the Bayes Factor, \( K \). For models \( M_1 \) and \( M_2 \), the Bayes Factor is given by:

\[
K = \frac{P(d | M_1)}{P(d | M_2)} = \frac{\int P(\theta_1 | M_1) P(d | \theta_1, M_1) d\theta_1}{\int P(\theta_2 | M_2) P(d | \theta_2, M_2) d\theta_2} = \frac{P(M_1 | d)}{P(M_2 | d)} \frac{P(M_2)}{P(M_1)},
\] (1.26)

where \( \theta_1 \) and \( \theta_2 \) are the best-fitting parameter-space vectors corresponding to models \( M_1 \) and \( M_2 \), respectively [47]. If \( K > 10 \), then model \( M_1 \) is strongly preferred over model \( M_2 \) [47]. Note that the integrals in Equation 1.26 are over each parameter, therefore the Bayes Factor will automatically penalise models with more parameters. In practice, the full Bayes Factor can be difficult to evaluate, in which case it can be estimated by statistics such as the Bayesian Information Criterion, as was done in this work.

1.7.1 Bayesian Information Criterion

The Bayesian information criterion (BIC) is an approximation of the more computationally complex Bayes Factor. The definition of the BIC is given in Equation 1.27, where \( d \) is the number of free parameters in the model and \( N \) is the number of data points:

\[
BIC = \chi^2 + d \ln \left( \frac{N}{2\pi} \right).
\] (1.27)

The BIC utilises the familiar \( \chi^2 \) statistic as a measure of fit quality but will also penalise the complexity of the model using the second term. For example, an arbitrarily complex model might yield a perfect fit to the data, but it would be penalised for having a high number of free parameters. Note that the value of BIC alone does not measure the goodness of a model. Rather, when comparing two
competing models, the change in BIC from the simpler to the more complex model is used as evidence in favour of or against the more complex model: if the inclusion of an additional parameter results in a decrease in BIC of over 10, the new model is strongly preferred [48, 49].

1.8 SPLASH

The SPLASH project (the Southern Parkes Large-Area Survey in Hydroxyl [1]) was designed to map the Southern Galactic Plane in all four ground-state OH transitions, addressing key issues on its global distribution, the degree to which it traces CO-dark molecular gas, and its role as a probe of density/temperature variations in the Galactic Disk.

This work provides a supplementary dataset to SPLASH, which has currently mapped Galactic longitudes $l = 332^\circ$ to $l = 10^\circ$ between Galactic latitudes $b = -2^\circ$ to $b = +2^\circ$. SPLASH detected OH both in emission and in absorption against the diffuse Galactic continuum background. Analysis of data from the pilot region yielded two observations that partially motivated this project:

1. As outlined in Section 1.4, OH is expected to trace $\text{H}_2$ in diffuse regions not traced by CO. Therefore SPLASH was expected to show OH detections without associated CO as has been seen in other studies [4, 22, 25, 50, 51]. This was not seen in the pilot region, and the reason why was unclear: Was the continuum brightness temperature ($T_c \approx 10$ K) too similar in value to the excitation temperature ($T_{ex} \approx 10 - 20$ K) thus preventing any measurable absorption\(^8\)? (We refer to this as the ‘contrast problem’.) Or is there some other reason why the OH gas is not detected?

2. The observations in the pilot region of SPLASH indicated that the OH gas was optically thin, which was expected from previous studies [52–54]. However, the Parkes beam is quite large ($\approx 12.5^\prime$), such that these low values of optical depth have a level of ambiguity: Is the gas actually optically thin, or is the beam just partially filled with optically thick clouds?

Our choice of the Australia Telescope Compact Array (ATCA; an interferometer), and of bright, compact background continuum sources allow us to answer both of these questions. The narrow beam of the ATCA in the 1.5D configuration used in this project (full width at half-maximum, FWHM $\approx 30''$)\(^9\) led to high background continuum brightness temperatures $T_c$\(^10\). These $T_c$ values

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\(^{8}\)See Equation 1.16. If $T_c \approx T_{ex}$, $T_b \to 0$.

\(^{9}\)Excluding antenna 6.

\(^{10}\)For a compact, unresolved source the measured brightness temperature scales as the inverse of the beam area.
were expected to be much higher ($\approx 1000$ K) than the excitation temperature of the OH gas ($5 - 15$ K), therefore avoiding the contrast problem suffered by SPLASH. High $T_c$ values (and the use of an interferometer) also enabled us to calculate optical depth directly, as described by Equation 1.24. This allowed us to confirm the low optical depths inferred from analysis of the SPLASH pilot region, and show that their result was not likely due to beam filling effects alone.

We now describe these observations (and the motivation behind our technical choices) in more detail in the following Chapter.
Observations and Data Reduction

2.1 Observations

High-sensitivity observations of OH absorption were carried out against a sample of 15 bright, compact continuum sources in the SPLASH survey region with the Australia Telescope Compact Array (ATCA). Note that much of the content of this section is based on the original ATCA project proposal for these observations (Project number C2976 [55]).

2.1.1 Seeking the ‘Missing’ Gas

As already mentioned above, one of the most surprising results from the SPLASH pilot region was the lack of evidence for OH outside the CO-bright regions of molecular clouds [1]. This was unexpected, given the high sensitivity of SPLASH, and the growing body of literature demonstrating the effectiveness of OH as a tracer of diffuse molecular gas. The authors believed that the similarity of OH excitation temperatures to the diffuse Galactic continuum background might be partially responsible (the ‘contrast problem’). These observations were designed to test this hypothesis and distinguish between cases where OH is present and unseen, and cases where it is genuinely absent (or of negligible abundance). The observations carried out for this work were designed to allow a $3\sigma$ detection sensitivity of $\tau \sim 0.01$–0.02 and hence can test whether non-detections in the Parkes data contain significant OH. Most sightlines are expected to contain at least one CO emission component in which no OH was detected in SPLASH (based on analysis of $\sim 25\%$ of the full survey region), meaning that any sufficiently bright continuum source in the survey region was a viable target.
2.1 Observations

2.1.2 Optical Depth Measurements

The observations carried out for this work were designed to measure $\tau$ in a large number of emission and absorption components throughout the inner Galactic Plane, providing a test of whether small optical depths are a general property of Galactic diffuse OH, or whether beam dilution masked the presence of higher $\tau$ material.

2.1.3 Source Selection

The 15 target continuum sources are shown in Fig 2.1. This sample size was chosen as a rough minimum required to obtain useful statistics on the optical depth distribution, and the potential impact of the contrast problem on OH detections in SPLASH. It was required that sources be bright, compact (unresolved or with sufficient unresolved structure at a beam size of $\sim 30''$), and located between $332^\circ < l < 8^\circ$, $|b| < 2.1^\circ$. The brightness temperature of an observed source is proportional to its spectral flux density (measured in janskys where $1\text{Jy} = 10^{26} \text{W m}^{-2} \text{Hz}^{-1}$) but inversely proportional to the area of the synthesised beam. Therefore we required sources with a spectral flux density $\sim 1\text{ Jy}$ at 1.6 GHz, which would result in brightness temperatures of $\sim 1000 \text{ K}$ when observed with our array configuration.

The flux density requirement also ensured that a $3\sigma$ sensitivity of $\tau \sim 0.01–0.02$ could be achieved in a reasonable integration time. Distant sources were preferable since they probe a larger number of absorbing components along the line of sight. However, the number of extragalactic and far-side Galactic sources fulfilling the criteria was small, and the target list therefore included several nearside HII regions with evidence for bright and compact substructure and intervening H\textsc{i} absorption. Our sources were selected from the 843 MHz Molongo Galactic Plane Survey catalogue (MGPS, [56]), and from the Southern Galactic Plane Survey (SGPS, [57]) and the National Radio Astronomy Observatory (NRAO) Very Large Array (VLA) Sky Survey (NVSS, [58]) 1.4 GHz continuum images. All sources were cross-checked against the recombination line measurements of Caswell (1987, [59]) in order to discriminate between HII regions and other source types, and were also examined for evidence of H\textsc{i} absorption in SGPS datacubes, in order to confirm nearside or farside Galactic distances, where relevant. 1.6 GHz flux density estimates were obtained estimating spectral indices from existing flux density measurements. For marginally resolved sources, it was assumed that at least 50% of the peak flux density would be lost as a result of our smaller beam size ($\sim 45''$ in MGPS and NVSS compared to $\sim 30''$ in this work).
2.2 Data Reduction

Figure 2.1: Target distribution overlaid on map of the SPLASH Phase 1 survey coverage. Black sources are nearside H\textsc{ii} regions, red sources are farside H\textsc{ii} regions, and blue sources are extragalactic. Large, medium and small crosses are sources with $S \gtrsim 2$ Jy, $S \approx 1$–2 Jy and $S \lesssim 1$ Jy. The source at Galactic longitude 349.73 was not imaged.

2.1.4 Observing Strategy

The CFB 1M-0.5k mode on the ATCA Compact Array Broadband Backend (CABB) was used to simultaneously observe all four ground state OH lines in zoom bands centred on the line rest frequencies. This provided a raw channel width of 0.09 km s$^{-1}$, which was binned to a width of 0.36 km s$^{-1}$ in this work, sufficient to resolve the narrowest OH features in SPLASH. The synthesised beam size needed to be small enough to achieve high brightness temperatures for compact sources, but large enough that the sources were not significantly resolved. The 1.5D array was therefore chosen, for a synthesised beam size of $\sim 30''$ at 1.6 GHz, which for a 1 Jy source corresponded to a brightness temperature of $\sim 500$ K. The project was awarded a total observing time of 50 hours for all 15 sources.

2.2 Data Reduction

The data gathered from the ATCA was reduced using the \textsc{miriad}\textsuperscript{1} package [60]. \textsc{miriad} is able to automatically flag the bad channels associated with known issues with the CABB system, such as the self-interference due to 640 MHz clock harmonics, lower sensitivity in edge channels, and channels known to contain radio frequency interference (RFI). The raw data for each background continuum source and the three calibrator sources (1934-638, 1827-360 and 1740-517) was split into individual frequency zoom bands centred on 1612 MHz, 1666 MHz and 1720 MHz. Any visibility ranges (channels or times) in the $uv$ files that were clear outliers were flagged as likely RFI. Calibration

\textsuperscript{1}http://www.atnf.csiro.au/computing/software/miriad/
2.2 Data Reduction

Figure 2.2: Positions of the antennas in the 1.5D configuration of the Australian Telescope Compact Array (ATCA). Antenna 6 is not included in this graphic as baselines including antenna 6 were not used in this project.

Figure 2.3: The top centre panel shows a sample of the $uv$-coverage from 12 hours’ observations using the ATCA in the 1.5D configuration (excluding antenna 6). The synthesised beam and the observed image from the compact model image at top left are shown. This figure was produced using the Friendly Virtual Radio Interferometer software (https://crpurcell.github.io/friendlyVRI/).
solutions were found using 1934-638 for flux and bandpass calibration, and 1827-360 and 1740-517 as phase calibrators (1827-360 for target sources at Galactic longitudes greater than 344°, and 1740-517 for the remaining target sources). Several of the antennas across the five days of observing demonstrated significant changes in phase between observations of the calibrator, so self-calibration was used to interpolate between these times using the data from the target sources. These calibration solutions were then applied to the visibility files, which were then inverted, cleaned and restored into continuum-subtracted brightness temperature data cubes and continuum images for the 15 sightlines and four OH ground state transition frequencies.

As described above, images of the candidate objects at the target frequencies and at the necessary resolution were not available prior to this work, so sources were chosen based on estimates or upper limits on their angular size. In total, 10 of the 15 sources were in fact found to be embedded in diffuse continuum emission or have significant large-scale structure that was poorly-sampled by the 1.5D array. This manifested as ‘waves’ or ‘ripples’ in the image-plane, rendering flux measurements inaccurate. We mitigated against this problem by flagging out short baselines corresponding to broad spatial scales. The minimum $uv$-distances used for each continuum source are given in Table 2.1. This removal of shorter spacings improved the continuum images but also reduced the signal to noise ratio.

Our final per-channel 1σ optical depth sensitivities are listed in Table 2.1. It can be seen that these are between 0.007–0.05, and are hence in general below the target sensitivities of the observations. If the original planned channel width of 0.7 km s$^{-1}$ is used, 80% of sightlines have $3\sigma_r \sim 0.015–0.055$. While this is lower than hoped, it is sufficient to carry out the aims of this project. We may also be able to refine the data reduction strategy further in the future to improve on this.
### Table 2.1: Detailed information for the continuum sources coinciding with the sightlines examined in this project.

<table>
<thead>
<tr>
<th>Source</th>
<th>$T_c$(1667) (K)</th>
<th>Type$^a$</th>
<th>Location in disk</th>
<th>$uν_{min}^b$ (kHz)</th>
<th>$σ_τ^c$</th>
</tr>
</thead>
<tbody>
<tr>
<td>G334.72+0.65</td>
<td>781</td>
<td>H II region (-16 km s$^{-1}$)</td>
<td>likely far-side</td>
<td>3</td>
<td>0.05</td>
</tr>
<tr>
<td>G336.49-1.48</td>
<td>3567</td>
<td>H II region (-25 km s$^{-1}$)</td>
<td>near-side</td>
<td>1</td>
<td>0.021</td>
</tr>
<tr>
<td>G340.79-1.02</td>
<td>1730</td>
<td>H II region core (-25 km s$^{-1}$)</td>
<td>near-side</td>
<td>1</td>
<td>0.017</td>
</tr>
<tr>
<td>G344.43+0.05</td>
<td>1384</td>
<td>H II region (-67 km s$^{-1}$)</td>
<td>near-side</td>
<td>0</td>
<td>0.026</td>
</tr>
<tr>
<td>G346.52+0.08</td>
<td>724</td>
<td>H II region (2 km s$^{-1}$)</td>
<td>likely far-side</td>
<td>2</td>
<td>0.03</td>
</tr>
<tr>
<td>G347.75-1.14</td>
<td>2277</td>
<td>VCS4 J1717-3948</td>
<td>extragalactic</td>
<td>0</td>
<td>0.007</td>
</tr>
<tr>
<td>G348.44+2.08</td>
<td>949</td>
<td>likely extragalactic</td>
<td></td>
<td>2</td>
<td>0.021</td>
</tr>
<tr>
<td>G350.50+0.96</td>
<td>1521</td>
<td>H II region (-10 km s$^{-1}$)</td>
<td>near-side</td>
<td>0</td>
<td>0.026</td>
</tr>
<tr>
<td>G351.56+0.20</td>
<td>1427</td>
<td>H II region (-42 km s$^{-1}$)</td>
<td>far-side</td>
<td>2</td>
<td>0.015</td>
</tr>
<tr>
<td>G351.61+0.17</td>
<td>1404</td>
<td>H II region (-42 km s$^{-1}$)</td>
<td>far-side</td>
<td>2</td>
<td>0.014</td>
</tr>
<tr>
<td>G353.41-0.37</td>
<td>1026</td>
<td>H II region (-16 km s$^{-1}$)</td>
<td>near-side</td>
<td>1</td>
<td>0.023</td>
</tr>
<tr>
<td>G356.91+0.08</td>
<td>790</td>
<td>Nonthermal</td>
<td>likely extragalactic</td>
<td>1</td>
<td>0.026</td>
</tr>
<tr>
<td>G003.74+0.64</td>
<td>598</td>
<td>likely extragalactic</td>
<td></td>
<td>0</td>
<td>0.022</td>
</tr>
<tr>
<td>G006.32+1.97</td>
<td>720</td>
<td>likely extragalactic</td>
<td></td>
<td>1</td>
<td>0.017</td>
</tr>
<tr>
<td>G007.47+0.06</td>
<td>525</td>
<td>H II region (-17 km s$^{-1}$)</td>
<td>far-side</td>
<td>0</td>
<td>0.035</td>
</tr>
</tbody>
</table>

$^a$ Velocities of radio recombination lines given for all H II regions.

$^b$ Background continuum sources that were found to have extended structure on the scale of the ATCA synthesised beam were reduced using only $uν$-spacings higher than this limit.

$^c$ Average 1σ optical depth sensitivity acheived along each sightline.
3

Gaussian Decomposition

The premise of Gaussian decomposition is that we should be able to take an observed spectrum and decompose it back into its individual Gaussian components. A given set of components represents a model of the clouds along the line of sight, which is evaluated by comparing it to the observed spectrum. If the model is a ‘good fit’ to the data, then the physical properties of the individual components are obtained from the model parameters. However, the fitting process can be complicated when the signal-to-noise ratio is poor and/or when multiple clouds at similar bulk velocities are blended along the line of sight.

The reliable decomposition of complex and blended spectra is an ongoing issue in many areas of astronomy, and as such this project aimed to devise a partially automated, repeatable method of Gaussian decomposition that removed as much subjective input as practical. The motivation for this aim is both the time-consuming nature of identifying Gaussian components manually, and the volume of spectra that are generated by modern surveys: it is becoming increasingly impractical to approach this problem without some degree of automation. Though the automated method described in this chapter still required a degree of subjective assessment, a significant amount of subjectivity and manual labour was removed.

3.1 Challenges of Gaussian Decomposition

In brief, the process of Gaussian decomposition involves the identification of a set of N Gaussian components in the spectrum, and the iteration over possible sets of parameters for those components to obtain a model that best fits the data. This iteration is guided by computing an evaluation statistic which is minimised in the case of the ‘best’ model\(^1\). From this brief introduction, the first challenge of Gaussian decomposition is revealed: *how complex a model is justified by the data?* If we wish

\(^1\)Thus the choice of evaluation statistic will necessarily have an impact on which model will be deemed to be the ‘best’.
3.2 Autonomous Gaussian Decomposition

to construct the ‘best’ model of a noisy data set (i.e. the model that has the lowest sum of squared residuals), we could naively keep adding Gaussian components until the residuals approach zero. However, this model would fit noise as well as ‘real’ features and would have little physical meaning. Thus we prefer an evaluation statistic that penalises arbitrarily complex models. This led us to evaluate our models using the Bayesian Information Criterion (BIC), as discussed in Section 1.7.1.

A second consideration arises from our need to develop a Bayesian method of Gaussian decomposition. A Bayesian approach requires us to compare different models to one another, rather than simply comparing them to the data. We therefore must begin with the simplest model (i.e. a flat spectrum), then add components to the model one by one. The addition of each component must lead to a significant improvement to the model, otherwise it is not included. Therein lies the second consideration: how can we identify the maximum number of reasonable Gaussian components contained in a spectrum in an automated and repeatable way? To do this we employ the autonomous Gaussian decomposition (AGD) algorithm GAUSSPY [61], which is described in Section 3.2.

A third and final consideration specific to the ground state transitions of OH is that the Gaussian models describing those four spectra along a single sightline should not be independent of one another. Though it is possible to see a spectral feature in one or more of the four line(s) without seeing it in the remaining line(s), a more typical profile would contain some sort of feature at the same velocity in all four transitions. We can use this information as a Bayesian prior when evaluating models. Many packages exist to achieve this (i.e. MPFIT [62]), but we opted to use the Python module EMCEE due to its versatility in applications throughout the scientific community. EMCEE is a Python implementation of the Markov Chain Monte Carlo (MCMC) technique [63]. The process of linking our model line parameters is discussed in Section 3.3.

3.2 Autonomous Gaussian Decomposition

The autonomous Gaussian decomposition (AGD) algorithm GAUSSPY identifies the locations of possible Gaussian components by finding peaks in the second derivative of the data\(^2\). Before finding the second derivative, GAUSSPY must first smooth the data using ‘regularised differentiation’ as described by Tikhonov (1963 [64]). This smoothing is achieved via a parameter ‘log(α)’ that represents a trade-off between smoothness and data fidelity [61]. A large value of log(α) (i.e. 2, 3) will lead to a high degree of smoothing resulting in the identification of fewer, wider Gaussian components. A small value of log(α) (i.e. -3, -2) will lead to a low level of smoothing resulting in the identification of

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\(^2\)This will identify emission. The spectrum is then multiplied by -1 and GAUSSPY is run again to identify absorption.
3.2 Autonomous Gaussian Decomposition

Figure 3.1: The panel at left shows the process by which the autonomous Gaussian decomposition (AGD) algorithm \textsc{gausspy} identifies the location of possible Gaussian components. A synthetic spectrum (black) is composed of three Gaussian components (green). The second derivative of the synthetic spectrum is shown (solid red line), and the $x$-values of the minima of the second derivative – the initial guesses for the locations of Gaussian components – are indicated on the synthetic spectrum (blue circles). The panels at right illustrate how a noisy spectrum must first be smoothed before differentiating. The top panel shows a noisy synthetic spectrum (grey), with an underlining model containing two Gaussian components (at $x = 25$ and $x = 75$, dashed line). The bottom panel shows the first derivative of the noisy data using a finite-difference-based numerical derivative (grey), along with regularised derivatives (i.e. ‘smoothed’) using the $\log(\alpha)$ values shown in the legend. These illustrations have been taken from Figures 1 and 2 from [61].

more, narrower Gaussian components. This is illustrated in Figure 3.1. After identifying the position of potential Gaussian components, \textsc{gausspy} will then optimise the parameters of those components via a least-squares minimisation, thus improving upon its initial guesses.

\textsc{gausspy} was written to analyse H\textsc{i} emission spectra from the 21-SPONGE survey [65], so the intended use of the $\log(\alpha)$ parameter was to identify H\textsc{i} features within a given range of widths: large values of $\log(\alpha)$ would identify the more broadened features of the warm neutral medium (WNM), while smaller values would identify the narrower features of the cold neutral medium (CNM). The WNM has a temperature of $\sim 10^3$–$10^4$ K while the CNM has a temperature of $\sim 10^2$ K [6]. Relative to the H\textsc{i} of the 21-Sponge survey, our data are expected to represent a much smaller range of kinetic temperatures ($\sim 10^1$ K). Further, the line widths of OH are expected to be determined more by turbulent rather than thermal broadening, so they were expected to be much narrower, and more similar to one another than those in the 21-Sponge survey. Therefore a single value of $\log(\alpha)$ was expected to be sufficient to identify the maximum number of reasonable components along a given line of sight,
while smaller values would identify noise peaks.

### 3.3 Gaussian Decomposition Method

The method we employed in this work is outlined in Figures 3.2 and 3.3.

The AGD algorithm gausspy [61] replaced the task of manually identifying potential Gaussian components. Its creators assert that gausspy identifies largely the same features as would be identified manually, so we employ it for its repeatability. An objective measure was needed to determine at what point the \( \log(\alpha) \) value began identifying noise peaks. We found that if we measured the Bayesian Information Criterion (BIC) of the models produced by each \( \log(\alpha) \) value, the model with the lowest BIC tended to identify all the components that would have been identified manually, along with additional components that either fit noise peaks or had very large \( \Delta v \) and low \( \tau_{\text{peak}} \). Thus we took the mean velocities of components identified in the model with the lowest BIC on for further analysis, as illustrated in green in Figure 3.2.

The lists of mean velocities for each ground state transition line were then compared so that components could be matched across the spectra, as illustrated in blue in Figure 3.2. Any Gaussian components identified within 1 km s\(^{-1}\) of one another across multiple lines were assumed to represent the same OH gas cloud. This limit was chosen based on the velocity resolution (0.36 km s\(^{-1}\) per channel) and the average feature width (full width at half-maximum \( \approx 3 \) km s\(^{-1}\) ), and is illustrated by the components labelled ‘\( v_1 \)’ in Figure 3.2. Based on preliminary modelling, any features separated by 1 km s\(^{-1}\) or less would not be able to be distinguished by our Gaussian decomposition method.

If a component was identified in only one of the four lines, it was also included (i.e. the component labelled ‘\( v_2 \)’ in Figure 3.2). Any Gaussian components within 4 km s\(^{-1}\) of one another in the same line were categorised as ‘blended’ (i.e. the components labelled ‘\( v_3 \)’ in Figure 3.2). This was based on preliminary modelling that showed that the detection rate of blended features separated by more than 4 km s\(^{-1}\) approached that of isolated features.

Any components that were not blended were categorised as ‘isolated’. This list of isolated and blended components for each line of sight was then taken to the next stage of analysis, illustrated in Figure 3.3.

A Bayesian approach to model selection requires the justification of any increase to the number of parameters in the model. All models were generated using the Python implementation of the Markov-Chain Monte-Carlo technique emcee [63], which allowed us to generate models that fit all four ground state lines simultaneously. Each model was built from components that had 6 parameters:
3.3 Gaussian Decomposition Method

Figure 3.2: Flow chart illustrating the process of Gaussian decomposition from the initial optical depth spectra to the identification of potential mean velocities of Gaussian components across the four ground state transitions of OH. The process continues in Figure 3.3.
Figure 3.3: Continued from Figure 3.2. Flow chart illustrating the process of Gaussian decomposition from the list of potential mean velocities of Gaussian components across the four ground state transitions of OH to the final model.
The values of \( v_0 \) were allowed to vary up to 1 km s\(^{-1} \) from their initial values. \( \Delta \nu \) and \( \tau_{\text{peak}} \) were limited to a ‘reasonable’ range of 0 – 12 and \( \pm 10 \), respectively\(^3\). The Markov chains were checked for convergence before the BIC of the model was calculated.

Our process of model selection began with the simplest model, the so-called ‘null-model’ (a flat spectrum), and a calculation of its BIC (BIC\(_0\)). To this model we then added a component at the location of a single isolated feature, following the flowchart shown in Figure 3.3. If this component reduced the BIC by at least 10, then it was included in the model (if not, it was rejected) and the next isolated component was trialled\(^4\). Once the list of isolated features was exhausted, the blended features were considered. First, a single component was placed at the velocity of the blended feature, the BIC of the resulting model was calculated (BIC\(_1\)) and compared to BIC\(_0\). If BIC\(_0\) – BIC\(_1\) > 10, then another component was added to the feature. The BIC of this two-component model was then calculated (BIC\(_2\)) and compared to that of the one-component model (BIC\(_1\)). If BIC\(_1\) – BIC\(_2\) > 10, then another component was added to the feature, and so on. If the addition of any further component failed to reduce the BIC by at least 10, then it was rejected and the previous set of components were added to the model. Once the list of blended features was exhausted, the model was complete.

In this method we used the presence of a feature in any of the four ground state lines to infer that a feature should likely be present in all of the lines. This is in contrast to a non-Bayesian approach that would only consider a single line at a time. This complicates the interpretation of the significance of our detections. However, for simplicity we show in Table 3.1 the number of components at the 1 and 2\( \sigma \) level of significance. Any components falling below the 1\( \sigma \) limit were not considered in further analysis, though any between 1 and 2\( \sigma \) were retained, pending further more robust statistical assessment.

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\(^3\)Ideally neither \( \Delta \nu \) nor \( \tau_{\text{peak}} \) would be restricted, but in practice it is useful to restrict the parameter space explored by emcee to increase its efficiency.

\(^4\)It is valid to compare the BIC of models containing single isolated components to that of the null model. The presence or absence of other components in the model do not affect the amount by which an isolated component changes the BIC, as it is a function of the \( \chi^2 \) statistic and a penalty based on the number of points in the data and the number of parameters added.
### Table 3.1: Significance of detections across the four $^2\Pi_{3/2}$, $J = 3/2$ ground state transitions of OH. Note that the main lines tended toward more significant detections than the satellite lines. Detections below the $1\sigma$ level were excluded from all further analysis.

<table>
<thead>
<tr>
<th>Transition (MHz)</th>
<th>Significance of Detections</th>
<th></th>
<th></th>
<th></th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>($\tau_{\text{peak}} &lt; 1\sigma_{\tau}$)</td>
<td>($1\sigma &lt; \tau_{\text{peak}} &lt; 2\sigma_{\tau}$)</td>
<td>($\tau_{\text{peak}} &gt; 2\sigma_{\tau}$)</td>
<td></td>
</tr>
<tr>
<td>1612</td>
<td>13</td>
<td>4</td>
<td>11</td>
<td></td>
</tr>
<tr>
<td>1665</td>
<td>4</td>
<td>8</td>
<td>16</td>
<td></td>
</tr>
<tr>
<td>1667</td>
<td>2</td>
<td>8</td>
<td>18</td>
<td></td>
</tr>
<tr>
<td>1720</td>
<td>11</td>
<td>7</td>
<td>10</td>
<td></td>
</tr>
<tr>
<td><strong>Total</strong></td>
<td><strong>31</strong></td>
<td><strong>27</strong></td>
<td><strong>55</strong></td>
<td></td>
</tr>
</tbody>
</table>

3.3 Gaussian Decomposition Method
Results and Discussion

4.1 Gaussian Decomposition Results

A total of 55 Gaussian components from 28 distinct OH clouds were identified across 10 of our 15 examined sightlines. A sample of the Gaussian models generated from these components are shown in Figure 4.2 which includes an example of a blended feature (G336.49-1.48) and a set of isolated features (G350.50+0.96). The parameters of the Gaussian components (mean velocity $v_0$, full width at half-maximum $\Delta v$, and peak optical depth $\tau_{\text{peak}}$ for each of the four OH lines) are shown in Table 4.1.

The threshold for detections used in this work was the requirement that the new model improve the BIC by at least 10. We found from preliminary modelling that for an isolated Gaussian feature (separated by more than 4 km s$^{-1}$) this corresponded to a significance of $1.3\sigma$. For blended features, the relationship was more complicated, following a rough power law $\sigma = 10.5x^{-1.5}$ where $\sigma$ is the signal to noise ratio required for a decrease of 10 in the BIC for a separation of $x$. The only blended feature that didn’t satisfy this limit was G351.61+0.17 at -42 km s$^{-1}$: this feature was judged by eye to be composed of two Gaussians based on the difference in center velocity of the features in the main lines.

Recall that one of our initial aims in this work was to resolve the ‘contrast problem’ suspected of possibly preventing the detection of significant amounts of OH gas in the SPLASH pilot region. If SPLASH failed to detect a number of OH clouds due to the lack of contrast between their excitation temperatures and the diffuse background continuum, then our narrow sightlines towards bright background continuum sources should have yielded significantly more OH detections. While we are yet to attempt a detailed component-by-component comparison, a preliminary inspection of the SPLASH spectra suggests that we did not detect any distinct OH clouds not seen in the SPLASH cubes.

Figure 4.1 reproduces Figure 5 from Dawson et al. (2014) [1], which illustrates how our ability to
Figure 4.1: Figure 5 from Dawson et al. (2014), showing the ranges of $T_{\text{ex}}$, $T_c$, and $N(\text{OH})$ for which OH would have been undetectable in SPLASH, assuming all continuum emission is located behind the OH gas. detect OH in that survey depends on $T_{\text{ex}}$, $T_c$ and $N(\text{OH})$. For $N(\text{OH})$ of $\sim 2 \times 10^{14}$ cm$^{-2}$, representative of that measured in this work (see Section 4.4 and Table 4.1), the excitation temperature and continuum background brightness must satisfy $|T_{\text{ex}} - T_c| \gtrsim 2$ K in order for SPLASH to detect the component [1]. The fact that we find no additional detections at this column density suggests that this condition tends to be fulfilled. This means that provided the column density is $\gtrsim 2 \times 10^{14}$ cm$^{-2}$, SPLASH is unlikely to miss gas due to lack of contrast between $T_{\text{ex}}$ and $T_c$. However, this does not rule out a scenario in which significant lower $N(\text{OH})$ material exists and is missed for this reason.

4.2 Optical Depth in the Galactic Disk

The sightlines in this work all pass through the Galactic Disk inside the solar circle, where most of the Milky Way’s molecular gas resides. The few previous studies that have measured OH optical depth in this region examined only a small number of sightlines with larger beam sizes than the present work (3 sightlines in [53] with a 3′ beam; 2 sightlines in [52] with an 3.5′ beam). While no firm conclusions can be drawn from these works alone, the measured values of $\tau_v$ were all small ($\lesssim 0.03$), consistent with our picture of optically thin OH throughout the Galactic Disk. We also compare our optical depth measurements to the work of Li et al. (2018 [4], henceforth referred to as LTN2018), who examined 44 sightlines toward bright extragalactic continuum sources using the Arecibo telescope. These sightlines were not confined to the disk, but covered the whole of Arecibo’s view of the sky. This presents us with the opportunity to compare the statistics of the optical depth distribution in the inner Galaxy with that in local ISM clouds. Figure 4.3 shows histograms of the optical depth detections made in this work, with comparison to those found by LTN2018 for the main lines.
The range of possible OH column density $N(\text{OH})$ values assuming the excitation temperature of the 1667 MHz line $T_{\text{ex}}(1667)$ lies between 5 and 15K.

The range of $\text{H}_2$ column densities using the $N(\text{OH})/N(\text{H}_2)$ factor of $10^{-7}$ from Nguyen et al. (2018 in preparation).
4.3 Excitation Temperature

Though at first glance the main-line optical depths of OH clouds detected in this work seen in Figure 4.3 appear higher than those found by LTN2018, our relatively high noise levels prevented the detection of lower values; our sample is incomplete below $\tau_{\text{peak}} \sim 0.06$. While it appears that the main-line optical depths also have a wider range than those seen in LTN2018, none are so high as to strongly invalidate the optically thin assumption. Indeed, $1 - e^{-\tau} \approx \tau$ within 10% for 93% of main-line detections in this work; therefore the assumption of low $\tau$ is likely to be appropriate for the interpretation of the SPLASH data set, which simplifies the calculation of $N($OH$)$. The satellite lines were found to have lower optical depth than the main lines with a roughly equal distribution to one another, though the low number of significant detections limits any further interpretation.

The 1667 MHz main line was seen in absorption in 26 of the 28 identified features, while the 1665 MHz line was seen in absorption in 24 features. The 1612 MHz line was $\sim 30\%$ more likely to be seen in emission (6 of 15 detections significant to $1\sigma$) when compared to the 1720 MHz line (5 of 17 detections). As outlined in Section 1.5.1, the interpretation of anomalous excitation in the satellite lines is complex, but we can broadly state that the conditions allowing emission in the 1720 MHz line to dominate that of the 1612 MHz line are not prevalent along the sightlines examined in this work. Modeling is now required to elaborate on these conclusions; there is a wealth of potential information in the excitation of the satellite lines that is often neglected in favour of the more straightforward main lines.

### 4.3 Excitation Temperature

As introduced in Section 1.3, excitation temperature $T_{\text{ex}}$ is a reparameterisation of the relative population of the levels involved in a given transition, and hence provides information about the excitation state of the gas. An estimate of $T_{\text{ex}}$ is also generally required to compute column density.

Here we explore two methods of obtaining excitation temperature from our observations. The first is the so-called ‘on-off’ method, employed extensively in measurements of HI emission/absorption, and described clearly by Anderson and Bania [68]. The second is a novel method recently outlined by Yan et al. [3], whereby the equation of radiative transfer (1.16) and the definitions of optical depth (1.9) and excitation temperature (1.4) can be uniquely solved for all four excitation temperatures and column densities, using measurements of the line brightness temperature $T_b$ and background continuum brightness temperature $T_c$ for all four OH lines. As promising as this sounds, we show in Section 4.3.2 that this method can not be applied to the observations made in this work, and suggest that it also can not be reliably applied more broadly.
4.3 Excitation Temperature

Figure 4.2: A sample of the Gaussian decomposition models obtained in this work for a blended feature (left) and a set of isolated features (right). The original data is shown in a lighter shade in each panel, and individual Gaussian components are shown in black. $^{12}$CO emission (from NANTEN [2]) is shown in grey. All other Gaussian models can be seen in Section B.

Figure 4.3: Optical depth histograms showing $\tau_{\text{peak}}$(1665) at left, $\tau_{\text{peak}}$(1667) at centre and $\tau_{\text{peak}}$ of the satellite lines (1612 and 1720 MHz) at right. The histograms of the main lines have been normalised and include results from LTN2018 [4] as a comparison. The dashed lines show the most extreme values of $2\sigma_{\tau}$ in the sample of detected sightlines, and give a rough illustration of the completeness limit of our sample.
4.3 Excitation Temperature

4.3.1 ‘On-Off’ method

The ‘on-off’ method takes advantage of the assumed arrangement of continuum and intervening gas cloud illustrated in Figure 1.6. If the OH gas cloud is assumed to be spatially larger on the sky than the bright, compact background continuum source, then observations ‘on’ and ‘off’ the background continuum source can allow us to find excitation temperature via the method that follows [36, 37, 43, 50, 52, 53].

Recalling the solution to the radiative transfer equation (1.16), in the case of our interferometric observations where \( T_c \gg T_{\text{ex}} \), the on-source (continuum-subtracted) brightness temperature is given by:

\[
T_{\text{ON}}^b = -T_c (1 - e^{-\tau_{\nu}}).
\]

From this we directly obtain \( \tau_{\nu} \). To compute \( T_{\text{ex}} \), we must then observe the cloud just ‘off’ the continuum source. The key assumption here is that the cloud is large and smooth compared to the size of the continuum, with constant optical depth across the region of interest. Off-source measurements are made with a single dish telescope to avoid resolving out real flux from the target cloud. The solution to the radiative transfer equation for the off-source position is:

\[
T_{\text{OFF}}^b = (T_{\text{ex}} - T_{\text{bg}})(1 - e^{-\tau_{\nu}}),
\]

where \( T_{\text{bg}} \) is the diffuse continuum background, consisting of any Galactic component plus the cosmic microwave background radiation (not present in Equation 4.2 since they are both negligibly small compared to \( T_c \), and also likely to be resolved out in the interferometric observations). Substituting \( \tau_{\nu} \) then allows us to solve for \( T_{\text{ex}} \), provided that \( T_{\text{bg}} \) is known.

In our implementation of this method, we used single pointings from the Southern Parkes Large-Area Survey in Hydroxyl (SPLASH [1]) to determine \( T_{\text{OFF}}^b \) and \( T_{\text{bg}} \). The Parkes beam FWHM is \( \sim 12.5' \) at 1666 MHz, and the effective resolution of the SPLASH data products is larger still, at \( \sim 15' \) [1]. This is almost 1000 times larger than the area of the ATCA beam and the compact continuum sources. A single pointing centred on the continuum source therefore represents an average off-source spectrum over a wide region surrounding the target position, and is expected to show little or no contribution from \( T_c \). Unfortunately, this approach was only possible for two of the sightlines in this work (G336.49-1.48 and G340.79-1.02) due to the limited availability of properly calibrated continuum data from SPLASH. However, it provides a useful test of the methodology.

For the purposes of this work, we computed \( T_{\text{ex}} \) on a channel-by-channel basis, rather than using our Gaussian models, as we do not always see a one-to-one correspondence between our Gaussian
components and the SPLASH spectra. This already hints that the assumption of an extended, smooth OH ‘cloud’ with constant properties across the SPLASH beam is likely not fulfilled. Nevertheless, we are able to obtain estimates of $T_{\text{ex}}$. Our derived excitation temperature spectra are shown in Figure 4.4. Focusing first on the main lines, we can see that in regions where no line is detected the spectra are noisy about the value of $T_{\text{bg}}$, then flatten within the ranges where lines are detected. While the spectra towards our two sightlines contain complex and blended features, we may obtain a crude estimate of a characteristic $T_{\text{ex}}$ by taking the mean over the FWHM of detected features. This results in values of $12 - 15$ K – consistent with the 5-15 K range expected.

While this agreement is encouraging, there are several additional caveats. Firstly, we have assumed that all of the diffuse continuum emission in $T_{\text{bg}}$ lies behind the OH clouds, which may not be the case in the inner Galaxy. Secondly, both sightlines show emission in one of the satellite lines in the ATCA data (1720 MHz in G336.49-1.48 at $-20$ km s$^{-1}$, 1612 MHz in G340.79-1.02 at $-27$ km s$^{-1}$), implying negative $\tau$ and hence negative excitation temperatures. However, this is not seen in the $T_{\text{ex}}$ spectra for the satellite lines, suggesting that all our results should be interpreted with extreme caution. In this context it is interesting to note that [52] successfully employed the same method to confirm negative excitation temperatures in the 1720 MHz line for two off-Plane OH clouds. Off-Plane sightlines are far simpler to analyse using this method, because the gas tends to be relatively local (therefore changing less rapidly on angular scales), there tend to be fewer blended components, and the continuum background is dominated by the CMB, and is therefore known to be behind the cloud. Further investigation is clearly needed into the reliability and limitations of this method in the complex inner Galactic Plane.

The potential importance of this modified ‘on-off’ method is illustrated by the upcoming GASKAP (Galactic Australian Square Kilometre Array Pathfinder) survey [69]. GASKAP will image the Galactic plane and Magellanic Clouds in 21 cm Hi emission and the 1612, 1665 and 1667 MHz OH lines at a spatial resolution similar to our ATCA observations. In sightlines towards compact bright background continuum where $T_{c} \gg T_{\text{ex}}$, GASKAP pointings can be used as ‘on’ positions where data from SPLASH (and its successors) can be used as ‘off’ positions. This opens the possibility of obtaining direct measurements of $T_{\text{ex}}$ over multiple sightlines across large regions of the sky. From this work it is not clear, however, whether these measurements would be sufficient to provide excitation temperatures with more precision than our broad estimate of 5-15 K. Our small sample size prevents further assessment of this method at present, aside from the fact that it serves as a useful verification of the expected excitation temperature range.
4.3 Excitation Temperature

4.3.2 Method to find excitation temperature numerically

As can be seen in Figure 1.2, the four $^2\Pi_{3/2}, J = 3/2$ ground state transitions of OH involve only 4 levels, so their excitation temperatures are not independent. Yan et al. [3] (hereafter YWD2017) attempt to exploit this fact to solve for the four column densities (and hence excitation temperatures and optical depths) from observations of $T_b$ and $T_c$ in all four lines. This method is motivated by the fact that when one combines Equations 1.16 with Equations 1.9 and 1.4 one can construct a set of four (non-linear) equations where the only unknowns are the column densities of the four levels. For each transition between upper level $u$ and lower level $l$:

$$T_b = \left( \frac{-h\nu_0/k}{\ln \frac{N_u g_l}{N_l g_u}} - T_c - T_{bg} \right) \left( 1 - e^{c^2 A_{ul} \left( N_u - \frac{g_u}{g_l} N_l \right) \phi_v} \right),$$  \hspace{1cm} (4.3)

where all symbols are defined as previously. YWD2017 assert that it is possible to then solve these equations to find a unique set of column densities and then excitation temperatures of all four lines.

In the case of this work where we have already measured optical depth directly from interferometric observations, we can construct an analogous system of equations by simply combining the expressions for $T_{ex}$ and $\tau_\nu$ (Equations 1.4 and 1.9) without referring to the solution to the radiative transfer equation\textsuperscript{2}. Combining Equations 1.4 and 1.9 to remove excitation temperature $T_{ex}$, we can construct a series of linear equations:

$$N_1 = \frac{5}{3} N_4 - f(1720),$$
$$N_2 = \frac{3}{5} N_3 - f(1612),$$
$$N_3 = N_1 + f(1667),$$
$$N_4 = N_2 + f(1665),$$  \hspace{1cm} (4.4)

where $N_1 - N_4$ are the column densities of the four OH ground state levels numbered from highest to lowest, and

$$f(\nu) = \frac{8\pi\nu^3}{c^3 A_{ul} \phi_v} \tau_\nu.$$

If we then attempt to solve these simultaneous equations, all the column density terms cancel, leaving only the expression

$$3f(1667) + 5f(1665) = 3f(1720) + 5f(1612),$$  \hspace{1cm} (4.5)

\textsuperscript{1}Where the number densities $n$ have been replaced with column densities $N$.

\textsuperscript{2}Even if we did explicitly include Equation 1.16 rather than solving it for $\tau_\nu$ first, our conclusions would remain the same.
4.3 Excitation Temperature

which reduces to the well-known optical depth sum-rule for the OH ground state lines \[70],

\[
\frac{\tau_{1667}}{9} + \frac{\tau_{1665}}{5} = \tau_{1720} + \tau_{1612}. \tag{4.6}
\]

The column densities in Equations 4.4 cancel due to the repetition of the statistical weights across the four lines, as these are what determine the coefficients of the column density terms; the only condition under which a unique solution could exist is if the four levels of the OH ground state had unique statistical weights. In fact, it appears likely that the full set of simultaneous equations given in Equation 4.3 also does not have a unique solution, contrary to the claims of YWD2017. This follows simply from the fact that to obtain the set of equations given in Equation 4.4, we simply expressed the relationship between \(\tau_\nu\) and \(T_{\text{ex}}\) in terms of the level populations. If these do not have a unique solution in their simplest form, then expressing the variables in terms of observables \((T_b, T_c\) and \(T_{bg}\)), as done by YWD2017 will not change this fundamental fact. The situation is complicated somewhat by the fact that YWD2017 did not actually present solutions to this set of equations; they incorporated both additional simplifying assumptions and additional constraints, based on factors unique to their own science goals regarding OH structure in the Galactic Centre. Therefore further examination is required before conclusions can be drawn regarding the broader applicability of their method.

Figure 4.4: Excitation temperature spectra for two sightlines within the SPLASH survey [1] using optical depth spectra obtained in this work. The shaded region in each plot shows the full width at half-maximum velocity extent of significant \(\tau_\nu\) measurements.
4.4 Deriving Column Densities

In the absence of further observational information constraining $T_{\text{ex}}$, a range of values from 5–15 K\(^3\) were used to determine a range of possible column densities along all sightlines [4, 22, 38, 50, 71–73]. This 5–15K range is also consistent with the results from the two sightlines for which the ‘on-off’ method was applied, and with the assumption that the main-line excitation temperatures should fall between the temperature of the local radiation field (minimum $\sim 3$ K from the cosmic microwave background) and the expected kinetic temperature of the gas ($\sim 15$ K). $N(\text{OH})$ was calculated using Equation 1.13 for the 1667 MHz line, and was found to be $\sim 1 - 40 \times 10^{14}$ cm\(^{-2}\) across the detected OH clouds. We convert these OH column densities to H\(_2\) column densities using the factor $N(\text{OH})/N(\text{H}_2) \sim 10^{-7}$ [74, 75] The results are tabulated in Table 4.1. Values of $N(\text{H}_2)$ calculated from $N(\text{OH})$ can potentially be compared to those calculated from CO to identify CO-dark molecular gas, even when distinct CO-dark components are not detected.

4.5 Detection of CO-dark OH Gas

As mentioned in Section 1.8, observations from the pilot region of SPLASH did not detect any OH regions without associated CO emission [1]. We compared our OH spectra towards high brightness temperature continuum sources to \(^{12}\text{CO}\) ($J=1\rightarrow0$) spectra from the NANTEN Galactic Plane Survey [2]. These spectra are plotted together in Figure 4.2 and in Section B.

From an initial visual inspection of the spectra, we detect 3 components with OH but possibly no associated CO in this work; their spectra are shown in Figure 4.5. Unsurprisingly, the amount of CO-dark molecular gas detected in this way will depend on the sensitivity of the CO data [76]. The CO data from NANTEN [2] used in this work have a rms noise of $\sim 0.25$ K in a 1.0 km s\(^{-1}\) channel.

We compare these results to those of Li et al. (2018 [4]; LTN2018) who identified 8 OH clouds with no associated CO out of 49 detected OH clouds in the Local ISM. Of these 8 clouds, only 3 were not blended with neighbouring features. The method by which they identified these regions involved Gaussian decomposition of main-line OH spectra from Arecibo and \(^{12}\text{CO}\) spectra from the Purple Mountain Observatory Delingha. Rather than applying a full Bayesian approach, they chose Gaussian components by eye and fit them via a least-squares minimisation technique. Given that their conclusions are in part based on the relative numbers of Gaussian components identified across their OH and CO spectra (and that many of the CO-dark OH clouds identified are blended), this method

\(^3\)Values resulting from $T_{\text{ex}}=1–20$K are shown in Tables A and A.
implies that further work is needed to confidently assert that our OH clouds are not accompanied by CO gas.

For the pilot region of SPLASH, Dawson et al. [1] used a different method: they compared OH and CO data cubes on a channel-by-channel basis and did not identify any channels with significant OH absorption but no significant CO emission. (More precisely, the number was consistent with statistical noise, although this does not preclude that some minimal fraction of detections were real.) This difference in methods may account for the lack of significant OH detections without associated CO in the SPLASH pilot region: they would only have been able to detect isolated CO-dark OH gas. Only one of the three possible CO-dark OH clouds detected in this work (out of a total of 28 clouds) was isolated, and only 3 of the 8 detected by LTN2018 (out of a total of 49 clouds) were isolated. Thus a full Bayesian approach to the Gaussian decomposition of both the SPLASH data and comparison CO data (similar to that outlined in Chapter 3) is required before any conclusions can be drawn.

Perhaps more significant is the possibility that the lack of CO-dark OH detections in the SPLASH pilot region was because it is not present in sufficient amounts for it to be detected. The expectation that SPLASH would detect significant OH not associated with CO was based largely on the detections of CO-dark OH clouds in the outer Galaxy [22] or in off-plane, local gas [25, 51]. Perhaps the conditions in the local ISM and outer Galaxy allow the existence of significant quantities of CO-dark OH gas, but those in the inner Plane do not. Indeed, the gas in the inner Galaxy is expected to have higher density and pressure, and contain more molecular gas than at the Solar radius, so it seems plausible that a higher fraction of the material should achieve sufficient densities and column densities to be CO-bright. There is observational support for this picture; in the local ISM, more massive molecular clouds are found to have lower CO-dark fractions [10], and the results of the GOT C+ survey suggest the the CO-dark fraction may decrease from ~50% at the Solar Circle, to as low as ~20% in the inner kiloparsec of the Galactic Plane [28].

Another relevant point that has not thus far been made is whether or not it is reasonable to expect to find isolated OH gas components without associated CO in surveys of the Galactic Plane. Through the work of Grenier et al. [10] we expect some diffuse molecular gas to be poorly traced by CO, and we assert along with others [1, 22–26] that OH will trace this gas. We also assume that the lack of CO in this gas is due to extinction effects: CO is more easily dissociated by ultraviolet radiation than H₂, implying that clouds will have a ‘surface layer’ of CO-dark molecular gas [21, 77, 78]. Therefore in order to detect OH without associated CO, one would presumably need to observe a sightline that slices through this surface layer of CO-dark molecular gas. The chances of this would depend on the physical extent of this CO-dark surface layer, the angular resolution of the observations and how many
clouds are expected to share the same bulk velocity and therefore ‘overlap’ spatially and spectrally. Figure 4.6 illustrates a toy model showing the effect of relative cloud size on the probability that observations along a given sightline will detect OH but not CO. It is expected that larger clouds will have a significantly lower percentage of CO-dark molecular gas than smaller clouds [10], therefore larger clouds may tend to have a thinner CO-dark surface layer. Figure 4.6 therefore highlights the need for detailed modeling of the expected cloud size and abundance, as well as the depth of the CO-dark layer in order to make meaningful predictions of CO-dark OH gas detections.
Figure 4.5: The three sightlines in this work suspected of showing CO-dark OH gas. The OH data is from this work, and the CO observations (in grey) are from the NANTEN telescope [2]. The mean velocities of OH features suspected of not having associated CO are indicated by a vertical blue line. Both the sightlines towards G336.49-1.48 and G340.79-1.02 show a blended feature in OH but seem to have only a single feature in CO. Based on the method used by LTN2018 these may be considered to be CO-dark. The sightline towards G351.56+0.20 shows a thin isolated feature in OH with no significant associated CO. The CO intensity scale is adjusted to fit the plot panels, and has no physical meaning.
Figure 4.6: Illustration of some of the factors affecting the probability that observations along a given sightline will detect OH but not CO. Ensembles of molecular clouds are shown in each panel, each with a core that is traced by CO and a surface layer that is not. It has been asserted [1, 22–26] that this surface layer of CO-dark molecular gas is traced by OH. The left-hand panel shows an ensemble of large clouds moving at the same bulk velocity and therefore not distinguishable by their spectra. The grid shows possible sightlines through this field, with those sightlines likely to detect OH but not CO highlighted in yellow. The right-hand panel shows the same but with smaller clouds.
Conclusions

In this work we set out the following aims:

- Develop a partially automated, repeatable method to identify OH containing clouds. This was achieved by the Bayesian Gaussian decomposition method described in Chapter 3.
- Determine the peak optical depth of those clouds in the four lines of OH. Of the 55 Gaussian components identified in this work, only two had main-line optical depths greater than 0.2. Thus we conclude that the OH gas in the Galactic plane is typically optically thin. This will simplify the calculation of $N$(OH) in the SPLASH data set.
- Test different methods of determining the excitation temperature of the OH clouds. The modified ‘on-off’ method described in Section 4.3.1 was only able to constrain the excitation temperatures of features along two sightlines to the already expected range of 5-15 K. The analytical method described by Yan et al. [3] was demonstrated to be invalid in the context of this work, and possibly invalid in principle.
- Determine the column densities of the OH clouds and use these to estimate the column density of H$_2$. OH column densities were estimated for a reasonable range of main-line excitation temperatures, resulting in values between $\sim 1 - 40 \times 10^{14}$ cm$^{-2}$. These were converted to H$_2$ column densities between $\sim 1 - 40 \times 10^{21}$ cm$^{-2}$.
- Identify whether any of these clouds were not associated with CO emission. Three OH clouds without associated CO emission were identified, though two of these were questionable as they formed part of a blended feature. The initial motivation of this aim was to resolve the $T_{\text{exc}}$-$T_c$ contrast issue identified in the SPLASH pilot region; our results suggest that SPLASH is unlikely to miss gas with $N$(OH)$\gtrsim 2 \times 10^{14}$ cm$^{-2}$, but leaves open the possibility that lower column density material may be undetected for this reason. Upon closer examination of methods used by others to identify CO-dark OH gas, a full Bayesian process of Gaussian decomposition
(perhaps similar to that outlined in Chapter 3) should be performed on both the SPLASH and CO observations before conclusions can be drawn.

This work has comprised a short study conducted over the period of the Masters of Research degree. In the months following, several aspects of the analysis will be refined, and the scientific implications explored in greater depth.

- We will revisit details of the data reduction process to attempt to improve upon the relatively high optical depth sensitivity limits in the data presented here. In particular, we will re-assess whether the impact of imaging artifacts from extended continuum structure on our $\tau$ can be quantified such that we can retain all baselines in our final data products, and whether this would improve the final signal-to-noise ratio.
- We will perform a better assessment of the detection statistics and optical depth completeness limits, so that our $\tau$ distribution may be more quantitatively compared with the literature on other regions.
- We plan to further investigate the validity of the method of determining excitation temperature described in Section 4.3.2, potentially by incorporating other constraints on $T_{\text{ex}}$ and $\tau$.
- We hope to further refine the ‘on-off’ method for determining $T_{\text{ex}}$ described in Section 4.3.1, through a closer examination of the key assumptions made in our application of the method. This includes exploring the effect of cloud structure on calculated excitation temperature (i.e. by not assuming a smooth cloud) and continuum source placement (i.e. by not assuming the continuum source lies behind the cloud).

Scientifically, we wish to pursue the following future goals:

- We will test whether our Bayesian Gaussian decomposition method can be applied to the full SPLASH data set, allowing us to perform much more thorough statistical analysis of the properties of OH clouds throughout the Galactic plane. This would allow meaningful comparisons with CO data, expanding our understanding of the properties and distribution of CO-dark OH gas.
- We will examine the relationship between OH optical depth and Galactocentric radius.
- We will investigate whether we can apply excitation and radiative transfer modelling to the SPLASH dataset, to constrain the properties of the OH gas from single-dish data alone.
- We may possibly adapt the work of large-scale galaxy disk simulations (e.g. [79, 80]) to perform extensive modeling to establish the likelihood of detecting CO-dark OH gas using various telescopes, thus informing future studies.
Column Density and Excitation Temperature
Table A.1: Values of OH column density $N$(OH) calculated using Equation 1.13 for a range of $T_{ex}(1665)$ ($^a$) and $T_{ex}(1667)$ ($^b$) values.
<table>
<thead>
<tr>
<th>Source</th>
<th>$v_0$ (km s(^{-1}))</th>
<th>$T_{\text{ex}}(1665)$ (K)</th>
<th>$T_{\text{ex}}(1612)$ (K)</th>
<th>$T_{\text{ex}}(1720)$ (K)</th>
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<td></td>
<td></td>
<td>1K</td>
<td>5K</td>
<td>10K</td>
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Table A.2: Values of excitation temperatures consistent with the values of column density calculated for a range of $T_{\text{ex}}(1667)$ values ($\theta$)
B

Gaussian Models
References


