Determining atomic hydrogen line parameters

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Level and stream: UG/PG (Science)
Level of difficulty: 2
Suggested readings: Thermodynamics; best fit to experimental data; Basics of Quantum mechanics, Hydrogen atom, Hyper-fine splitting in Hydrogen

AIM: Determination of HI line frequency, velocity width and temperature of emitting gas from the supplied data.

Software: anyone of the following:(i) GNU software to fit and plot data files: 'gnuplot', (ii) Generic data analysis software, e.g., Scipy, Octave or Matlab.

1 Introduction

Most of the mass of the interstellar medium is in the form of atomic hydrogen (HI).
The 21-cm line, for which the exact frequency is 1420.4057 MHz, corresponds to the transition between the two hyper-fine sub-levels of HI in ground state.

When the spins of the electron and of the proton are parallel compared to when they are anti-parallel, the energy differs by $6x10^{-6}$ ev. The parallel state has higher energy and is due to interaction of magnetic fields produced by the spinning electron and proton. The energy difference could be determined using quantum mechanics.

The excitation temperature ($T_{\rm ex}$) of the line is typically called Spin temperature T_s.
The transition is strongly forbidden, the spontaneous emission probability, $A_{\rm ul}$ = $2.87x10^{-15}$ s^{-1}, implying a radiative lifetime of the upper sub-level ($1/A_{\rm ul}$)=$1.1x10^7$ years.
This timescale is much larger than the time between H atom collisions at high densities on Earth. So, it cannot be observed in laboratory.

To determine the relationship of intensity of observed HI with the intrinsic properties of the emitting cloud, radiative transfer equation needs to be used.

$$\frac{\mathrm{dI}(\nu)}{\mathrm{ds}} = j(\nu) - k(\nu) I(\nu) \tag{1}$$

where, $I(\nu)$ is the specific intensity at frequency ν, s is the distance from observer along the line of sight, $k(\nu)$ and $j(\nu)$ are the volume absorption coefficients and emissivity respectively.

Rayleigh-Jeans approximation is valid at 21 cm, such that

we can write $I_\nu = 2kT_B\nu^2/c^2$, where T_B is the brightness temperature. Equation (1) can be written as:

$$\frac{dT_B(\nu)}{d\tau(\nu)} = T_s - T_B(\nu)$$

where, T_s is called the excitation or "spin" temperature.

Consider an HI cloud with atoms having number density n_2 in the upper level and n_1 in the lower level. Let the corresponding statistical weights be g_2 and g_1 respectively. Then, T_s is defined by, $\frac{n_2/g_2}{n_1/g_1} = e^{-E/kT_s}$.

When HI gas is in thermodynamic equilibrium at temperature T, T_s= T.

For the simple case of isolated, single homogeneous cloud, the above can be solved to yield,

$$\Delta T_B(v) = (T_s - T_{bg})(1 - e^{-\tau(v)}) \quad (2)$$

Here $\tau(v)$ is the optical depth at velocity v and is related to column density ($N(v)$) by

$$\tau(v) = \frac{N(v)}{C.T_s}$$

The constant (C) has a value of $1.83 \times 10^{18}\,\mathrm{cm}^{-2}K^{-1}(\mathrm{km}/s)^{-1}$.

When optical depth of the line is small, we can write (neglecting any background radiation [$T_{bg} = 0$])

$$T_B(v) = T_s.\tau(v) = N(v)/C \quad (3)$$

When the emission is optically thick, T_B = T_s, or its brightness temperature is the spin temperature.

HI natural width is extremely narrow as the lifetime of the upper level is only limited by collisions which are rare in the general ISM. Therefore, HI broadening is only caused by the Doppler effect.

For simple line profile, and optically thin emission, we finally get

$$N(\mathrm{HI}) \simeq 1.83x10^{18}\Delta T_B(v)\,\Delta v\ \mathrm{atom\,cm}^{-2} \quad (4)$$

where, Δv is the line full width at half maximum (FWHM) in km s^{-1}.

For large optical depth, we must use the complete expression.

For observations in the Galaxy, or in nearby galaxies, the radial velocity is generally given with respect to the local standard of rest (LSR).

1.1 Absorption Measurements

The 21-cm line can be seen in absorption in front of a continuum radio-source or in front of the 21-cm emission of warmer HI.

Assuming an uniform cloud and the continuum source intensity variation across the line to be negligible, we get

$$T_B(\nu) = T(1 - \exp[-\tau(\nu)]) + T_c \exp[-\tau(\nu)] \tag{5}$$

where T_c is the brightness temperature of the continuum.

Typically, absorption measurements are made with interferometers with a high angular resolution, which resolves out the extended emission from the cloud. In such cases, we get

$$T_{\mathrm{abs}}(\nu) = T_c.\exp[-\tau(\nu)] \tag{6}$$

Note that τ is inversely proportional to the excitation temperature (kinetic temperature), so that the column density of HI cannot be derived from an absorption observation alone.

If we observe the 21-cm emission in the directions immediately adjacent to that of a source with a large single antenna, by interpolation we obtain the expected **emission** that would be observed in the direction of the source in the absence of this source.

This provides

$$\Delta T_{\mathrm{em}}(\nu) = T(1 - \exp[-\tau(\nu)]) \tag{7}$$

From the above two equations (6, 7), we can, in principle, derive the kinetic temperature and the optical depth, hence the column density in the cloud.

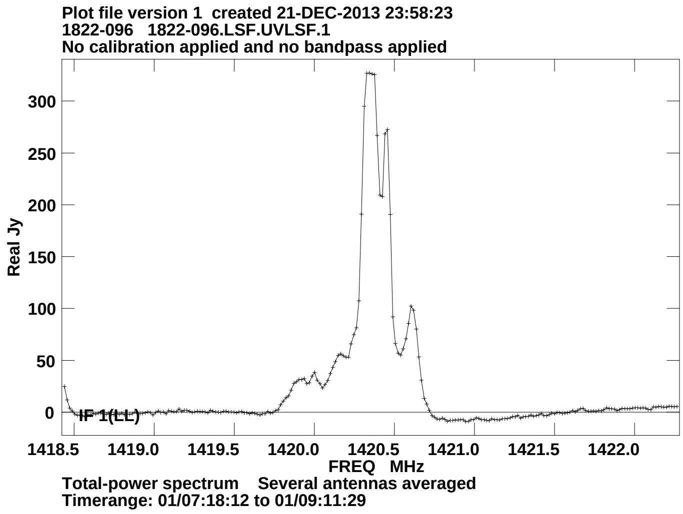

Figure 1. The 21-cm emission spectrum (Top) in the vicinity of the continuum source, calibrator 1822-096.

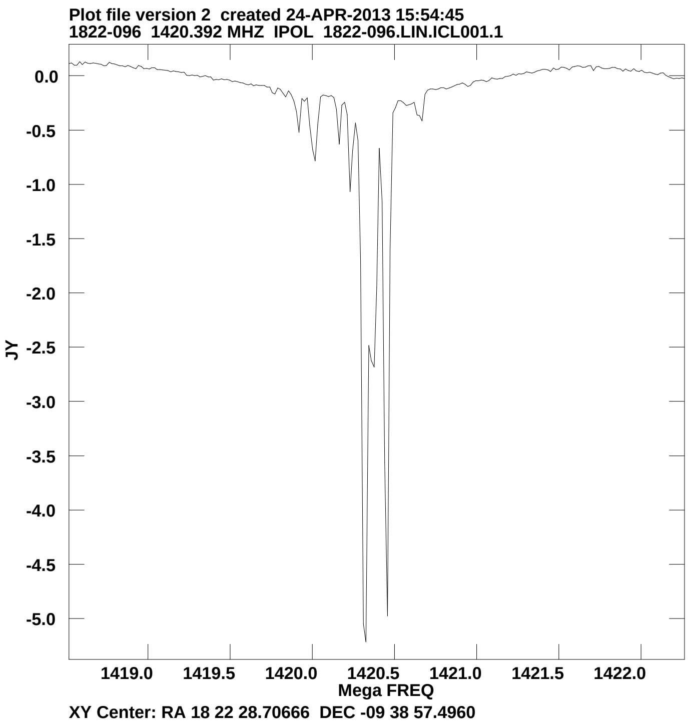

Figure 2. Absorption towards the same source as in Fig. 1 [both the observations were carried out with GMRT]

Application:

The main application of the HI line is the measurement of the mass, the distribution and the kinematics of atomic gas in our Galaxy and in external galaxies.

For this, it is generally assumed that the line is optically thin, so that the column density does not depend upon the physical temperature of the gas. This can be misleading as there is in general way to determine the optical depth.

So, the masses obtained should be considered as lower limit.

The atomic ISM is extremely inhomogeneous. The 21-cm emission is dominated by filaments, sheets and shells. This structure has a fractal appearance and could be due to turbulence.

There are 2 stable phases in the atomic ISM. One is warm (~3000-6000 K), low density (0.1-0.3 cm^{-3}), and is rarely seen in absorption (absorption coefficient is very small due to high T_s and low density).
The other one is cold (~60-100 K), denser (~10 cm^{-2}), and dominate absorption.
It can also be found in the envelopes of molecular clouds.
The warm component contains about as much matter as the cold component in the Galaxy, but it forms a thicker disk. Their respective mean half-thicknesses $|z|$ are about 186 and 105 pc near the Sun.

There exist some neutral gas in the halo, with a high velocity dispersion (60 km.s^{-1}) and a scale height of ~4 kpc.

HI is also found at high galactic latitudes that falls onto the galactic plane with velocities ranging from a few km.s^{-1} to several hundreds of km.s^{-1}.

These high-velocity clouds might be of extra-galactic origin (satellite dwarf galaxies), or originate from the hot ionized gas ejected by supernovae and chimneys from the galactic disk which then falls back onto the disk while cooling and recombining (Galactic fountain).

2 Experiment: Find the HI line parameters and rms

(i) See Figure 1, and 2 which shows HI emission and absorption towards the compact background source 1822-096.

Assuming the spectrum is composed of a few lines (each is of Gaussian shape on a baseline) find the frequency and width of them from both emission and absorption ?
Two of the associated data files are: (A) 1822-096.H1.abs (for absorption), and (B) 1822-096.self.spec (for emission).
Do the emission and absorption line widths differ greatly ?
If so, why ?

(ii) Using the line widths, determine the temperature of the emitting/absorbing gas responsible for the 2 main lines (most intense) in emission and absorption. Assume the line widths are due to Doppler motion (due to kinetic temperature) of the emitting atoms.

(iii) Using the absorption spectrum, find the rms noise (off from absorption).

(iv) There is another associated data file (C) '1822-096.offset.H1.abs', which has the spectrum of a region determined from a region away from the compact source.
Determine the rms across the band.
Note: In practice, a spectrum may contain certain offsets. Therefore, to find rms noise, you could first subtract any significant baseline (part without any visible line emission) offsets by either a constant term with a linear slope or a quadratic term, before determining rms noise from a few channels across the band.

(v) The calibrated *uv* (spatial frequency domain) data for each frequency channel which were used for making the spectrum, can be considered as a Fourier transform of the Compact source towards which the observations were made.
Then find the rms noise using (a) a single baseline (1-16) with 30 sec of data averaged (1822-096.spl.1-16.30s.spec), (b) the same baseline with 5m of data averaged (1822-096.spl.1-16.5m.data.spec) and (c) using data from the whole array (file: 1822-096.spl.full.data.vector.aver.spec)
(d) How do the above 3 cases compare with what you learnt in the school ?

References:

(i) Galactic & Extragalactic Radio Astronomy, by G. L. Verschuur & K. I. Kellermann.
(ii) The Feynman Lectures on Physics by Feynman, Leighton & Sands.
(iii) Data Reduction and Error Analysis for the Physical Sciences by P. Bevington & D. K. Robinson.