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I. Introduction

Discuss general viewpoint: What can happen to an atom or molecule or dust grain sitting in the ISM?

- Can it absorb a photon? Energy levels $\leftrightarrow h\nu$ (Need cross-sections for dust grains)
- Can it collide with other particles? $\rightarrow$ collisional rates [s^{-1} cm^{-3}]
- Is there a magnetic field? Is the particle charged?
- Are there cosmic rays? These can penetrate dense gas.

Three possible sources of ionization (and excitation):

a) photons b) collisions c) cosmic rays

Typical collisional energies: kinetic energy. Translate $mv^2 \rightarrow kT \rightarrow h\nu \rightarrow eV$.

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Dust grains also occur in the neutral medium, and probably also in the (warm) ionized medium. Dust grains play an important role in various processes:

- extinction of starlight
- emission of absorbed energy in FIR
- formation of molecules often occurs on grain surfaces
- absorption of ionizing UV radiation and Lyα photons (reducing amount of ionizing radiation)
- heating of HI gas by *photoelectric* emission

Composition: carbon and silicates. Typical sizes: 0.01 - 0.1 μm (How do we know? $\rightarrow$ shape of extinction curve). Grains as small as $\sim$ 60 atoms across discovered; evidence from emission lines in NIR and excess emission at 5 - 40 μm over what is expected from dust in the ISM. The larger dust grains have temperatures between 10 and 40 K, while the small ones can be heated to higher temps due to the absorption of even a single photon (smaller heat capacity, as volume $\propto r^3$). A promising candidate for small dust grains: polycyclic aromatic hydrocarbons ($\sim$ car soot!)

In hot environments, dust grains may be destroyed by *sputtering*, where collisiosn of grains
-9- with other atoms, electrons, or molecules knock molecules off the grains. At low temperatures, molecules stick to dust grains, causing depletion of heavy elements along certain lines of sight (most dust in the plane). Dust contributes about 1% of the mass of the ISM in the solar neighborhood, mostly in the form of large grains.

Ionized gas

a) Photoionization: especially effective near hot stars. Shock ionization
b) Cosmic rays: can occur throughout most of the ISM, so can also produce a small amount of ionization in denser gas (though recombination happens quickly, so not much of gas is in an ionized state at any given time).

-11- c) Collisions Hot ($\sim 10^7$ K) expanding bubble sweeps up shell of warm ($\sim 10^4$ K) ionized gas, which shows a different optical spectrum than HII regions. The hot gas shows up by:

- Free-free and line x-rays
- absorption lines of highly ionized species toward bright UV sources

Kirchoff's law: apparently can still see absorption lines when looking through a gas that is hotter than the source behind it!

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Magnetic fields and cosmic rays In the solar neighborhood, $\mathbf{B} \sim 2 - 5 \times 10^6$ Gauss. This follows from measurements of *Faraday rotation*, giving $< n_e B_{||} >$ toward pulsars and radio sources. The random component of the B field is probably as large as the uniform component. In clouds, the B field can be much higher, $\sim 70\mu$G (from Zeeman effect splitting measurements).

The magnetic field is important for several reasons:

1. It aligns elongated grains, giving rise to polarization of starlight
2. Causes relativisitic electrons to emit synchrotron radiation, and most likely plays a role in accelerating electrons to relativistic velocities ("magnetic bottle", Fermi acceleration).
3. Provides pressure support against gravitational collapse of matter since it is frozen into the matter due to ionization heavy elements. It also seems to play animportant role in solving the angular momentum problem in star formation.

Shu et al.

Total energy density of cosmic rays in solar neighborhood: $U_R \sim 1.3 \times 10^{-12}$ erg cm^{-3}. Why
-14- are cosmic rays important?

- They produce γ-rays through collisions with atoms and molecules. The observed γ-ray intensity from the ISM forms an excellent independent measure of the total amount of matter between stars. For example, calibrating the conversion factor of CO line intensity to H_2 mass.
- Provide pressure against gravitational collapse

Five pressures that play an important role in supporting the ISM against gravitational collapse:

1. thermal $P = nkT$
2. magnetic $P = \dfrac{B^2}{8\pi}$
3. turbulent (bulk motion)*
4. cosmic ray
5. radiation

* The cloud to cloud velocity dispersion due to turbulence on various scales increases line widths [over?] thermal widths. From the table on page -12-, you will note rough thermal pressure euqilibrium between the components. This is not a coincidence; in fact, some of this information was inferred by assuming pressure equilibrium. The argument is that if there

were no equilibrium, the resulting perturbations would be wiped out on sound-crossing time
-15- scales, which are short compared to the time scales we wouuld consider the ISM to evolve over.

However, the actual evidence for equilibrium in the thermal pressure is scarse, and there are claims that it is not true in the very local ISM[1].

There seems to be a “cosmic conspiracy”: the estimates for the thermal, magnetic, and cosmic ray pressure for the solar neighborhood give roughly equal numbers for all three. Thus it may be inappropriate to only consider the thermal pressure (the only one that can be measured with much certainty). Interestingly, the magnetic pressure number is also very similar to the energy density of the CMB. [2] [3]

-16- Next section is sort of a shortened condensation of Draine’s chapters 2 and 3. We may come back to specific topics discussed there in more detail.

II. Validity of the laws of statistical physics in ISM conditions

Four major laws of statistical physics:

1. **Maxwellian** velocity distribution
2. **Boltzmann distribution** of energy levels in atoms and molecules
3. **Saha equation** for ionization equilibrium
4. **Planck function** for radiation

Maxwellian T is defined by motion of particles. ($\vec{\omega}$ = velocity = $\vec{v}$ in Draine). $f(\vec{\omega})\mathrm{d}\vec{\omega}$ = fractional number of particles whose velocity lies within the three-dimensional volume element $\mathrm{d}\vec{\omega} = \mathrm{d}\omega_x \mathrm{d}\omega_y \mathrm{d}\omega_z$, centered at velocity $\vec{\omega}$.

In thermodynamic equilibrium, $f(\vec{\omega})$ is isotropic, so $\vec{\omega} \rightarrow \omega$.

$$f(\omega) = \frac{\ell^{3/2}}{\pi^{3/2}} \exp(-\ell^2\omega^2); \quad \ell^2 = \frac{m}{2kT} = \frac{3}{2 < \omega^2 >}$$

$$f(\omega) = \left(\frac{m}{2\pi kT}\right)^{3/2} \exp\left(-\frac{m\omega^2}{2kT}\right)$$

For two groups of particles with different masses, we replace ω by u, the relative velocity
-17- between the two groups, and m by the reduced mass $m_r = \dfrac{m_1 m_2}{m_1 + m_2}$. For H atoms colliding with particles of mass Am_H, Spitzer derives:

$$< u >= \left[\frac{8kT}{\pi m_r}\right]^{1/2} = 1.46 \times 10^4 \sqrt{T} \left(1 + \frac{1}{A}\right)^{1/2} \;\; [\mathrm{cm\ s^{-1}}]$$

…Speed vs. velocity

The Maxwell velocity distribution is characterized by several speeds:

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[1] Bowyer et al. *Nature* 1905

[2] Draine discusses possible reasons in section 1.3

[3] Supplemental info in Draine chapter 1.

- Most probable speed: $\omega_o = \sqrt{\dfrac{2kT}{m}}$

- RMS speed: $< \omega^2 >^{1/2} = \sqrt{\dfrac{3kT}{m}}$

- RMS velocity in one direction: $< \omega_x^2 >^{1/2} = \sqrt{\dfrac{kT}{m}}$

-18- **Boltzmann distribution** gives the population of energy levels in an atom or molecule:

$$\frac{n_u}{n_l} = \frac{g_u}{g_l} \exp\left[-\left(E_u - E_l\right)/kT\right]$$

where $n_{u,l}$ are the number densities, $g_{u,l}$ are the statistical weights, and $E_{u,l}$ are the energies of the levels. (Partition function: summing over all energy levels. . . chemical term).

Saha equation describes ionization equilibrium:

$$\frac{n_{i+1}}{n_i} = \frac{g_{i+1} g_e}{g_i} \left(\frac{2\pi m_e kT}{h^2}\right)^{3/2} \exp\left(-I/kt\right)$$

where I is the ionization potential for an ion in the ground state and initial ionization state i (aka, the energy required to ionize from i to $i+1$). $g_e = 2$ (two spin conditions). Electrons affect whether and how easily atoms can be ionized.

Plank function specifies the radiation field: [4]

$$\begin{aligned} B(\nu) &= \frac{2h\nu^3}{c^2} \frac{1}{\exp[h\nu/kT] - 1} \\ &= \sim \frac{2\nu^2}{c^2} kT \quad \text{for } h\nu << kT \text{ (Rayleigh} - \text{Jeans)} \\ &= \sim \frac{2h\nu^3}{c^2} \exp[-h\nu/kT] \quad \text{for } h\nu >> kT \text{(Wien)} \end{aligned}$$

III. Statistical equilibrium

-19- The four laws discussed above hold under thermodynamic equilibrium (TE). However, this is not often the case for the ISM. Thermodynamic equilibrium requires **detailed balancing**, i.e. each process is as likely to occur as its inverse. Example: Consider the 3727 Å emission from O^+. This is a forbidden transition (actually a doublet). The excitation of the electron level occurs through collisions with electrons, in most conditions in the ISM. If detailed balancing were to hold, de-excitation should also occur by collisions. However, as we will see, under the low density conditions found in the ISM, collisions are rare, and de-excitation is

[4] $e^x \approx 1 + x$ for $x << 1$
$e^{h\nu/kT} - 1 \approx 1$ for $h\nu << kT$
$e^{h\nu/kT} - 1 \approx e^{h\nu/kT}$ for $h\nu >> kT$

more likely to proceed through emission of a photon, in spite of the fact that we are dealing with a forbidden transition. Thus [OII] emission can be quite strong, and by converting collisional (kinetic) energy into radiation, we actually have created a cooling mechanism for the gas.

Another reason why TE does not hold is the strong **dilution** of the radiation field: the concept of a dilute radiation field is quite familiar. For example, the sun's photosphere is $\sim$ 6000 K, and at the surface the flux leaving the sun is approximately that of a blackbody of this temperature. However, the Earth is not 6000 K because by the time the radiation reaches us, it is diluted. A diluted radiation field is one in which the energy density does not match the color temperature.

-20- For the solar neighborhood, the total energy density of the radiation field due to all stars in that volume is about 1 eV cm^{-3} (close to cosmic ray density, as mentioned before). When interpreted as an average temperature using the Stefan-Boltzmann law (energy density of a blackbody, u = aT^4), this energy density implies an equivalent temperature of $\sim$ 3 K. Yet the color temperature implied by the shape of the spectrum of this Interstellar Radiation Field (ISRF) is that of A and B stars (T $\sim$ 10^4 K). So there is a **dilution factor** W given by:

$$W \approx (\frac{3}{10^4})^4 \approx 0.25 \times 10^{-14}$$

We conclude that using the Plank law to describe intensities is not correct.

What about the other laws?

1. Maxwell velocity distribution Good news! It is generally valid. Detailed balancing is possible for the elastic collisions that are generally occurring. Because the maxwellian distribution is a good description of the motions of the particles, we can define a kinetic temperature which describes the physical condition of the gas. Often, for a plasma, the kinetic temperature is equal to the electron temperature: $T_{ions} = T_e$. $T_{ions} \neq T_e$ may occur behind shocks.

-21- **2. Boltzmann distribution** is rarely correct. If excitation and de-excitation occurs by photons, we may still not have a Boltzmann distribution because the photon distribution is not given by the Plank function. Often we do not even have detailed balancing. However, as we will see, sometimes the distribution of excited levels is not too different from Boltzmann distribution. This happens when collisions dominate excitation and de-excitation, while radiation is relatively unimportant.

To describe situations close to TE, Spitzer introduced the so-called b-factors (Draine calls them "departure coefficients").

$$b_j \equiv \frac{n_j(\text{true distribution}}{n_j(\text{LTE distribution})}$$

Example: in an HII region, the hightest excited levels of HI have $b_j \sim 1$. Some radiation does escape (producing radio recombination lines) but collisions dominate the level populations. Since motions of particles *are* described by a Maxwellian velocity distribution, whenever collisions dominate the level population they will closely follow a Boltzmann law.

-22- In general, a Maxwellian velocity distribution tends to set up a Boltzmann population for energy levels in the atoms/particles *if* transitions resulting from emission and absorption of photons are relatively unimportant, and collisional (de-)excitation is dominant. In the case of the highly excited levels in H mentioned before, collisions with electrons are dominant.

3. Saha equation is generally not valid. There is no detailed balancing, and even though the ionization and recombination processes are each other's inverse, the ionization process is determined by the photon field in most cases, while the recombination process is determined by collisions between A^+ and e^-. The collision rate depends on $\{n_e, n_{A^+}\}$ and T_e, but the ionization is dependent on $T_{radiation} (\neq T_e)$.

So in general, assume **statistical equilibrium**, where there is a balance between transitions one way and the other way, no matter what process caused each transition.

In level i, we have n_i atoms cm^{-3}, and R_{ij} is the rate coefficient such that

$n_i R_{ij}$ [s^{-1}] = # transitions from level i to level j
$R_{ji} n_j$ [s^{-1}] = # transitions from level j to level i

$$\frac{\mathrm{d}n_i}{\mathrm{d}t} = \sum_j \left(-R_{ij} n_i + R_{ji} n_j\right); \quad i = 1, 2, \ldots$$

-23- In statistical equilibrium, $\frac{\mathrm{d}n_i}{\mathrm{d}t} = 0$. The rate factor R_{ij} includes *all* possible processes that would take the atom or molecule from level i to j. In the worst case, you would have to include many processes to calculate the n_i values. This requires knowledge of a lot of physical input parameters, e.g. cross sections for particular processes, collisional rate coefficients [5], etc.

In other cases, where only one or two processes matter, the situation can be very simple. We will encounter cases of each.

Before going more into Ch 2 and 3 in Draine, we will first discuss some basic radiative transfer.[6]

IV. Radiative Transfer

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Flux at surface of a sphere:

-25- $F_\nu = \pi B_\nu$ for blackbody

$F_\nu = \pi I_\nu$ for isotropic emitting non-blackbody

Flux at a distance r:

$F_\nu(r) = \pi I_\nu (\frac{R}{r})^2 = \frac{L_\nu}{4\pi r^2}$
where R = radius of body and L_ν = luminosity of body [erg s^{-1} Hz^{-1}]

Energy density of radiation

$u_\nu = \frac{1}{c} \int I_\nu \mathrm{d}\Omega$ [erg cm^{-3} Hz^{-1}]

[5] Drane section 2.1
[6] RL Ch 1, Draine Ch 6, 7

Radiation pressure

$P_\nu = \frac{1}{c}\int I_\nu \cos^2\theta \mathrm{d}\Omega$

Emission and absorption coefficients

V. Radiative transfer equation -26-

VI. Einstein coefficients -28-

VII. Line profile function, $\phi(\nu)$

-31- See RL chapter 10.6 and Draine 6.4

Natural line width

Key point: A small Einstein coefficient A results in a *narrow* line.

The natural line width of most transitions is quite small, and broadening due to other effects is more important.

Doppler broadening

- Thermal velocities
- Bulk motion (turbulence)

Collisional broadening

$\sim$ Pressure broadening, which is not generally important in the ISM because the density is so low…mostly occurs in stellar atmospheres. This still produces a Lorenzian profile, but with:

$$\phi(\nu) = \frac{4\Gamma^2}{16\pi^2(v - v_o)^2 + \Gamma^2}$$

VIII. Atomic H in the ISM

-33- Draine Ch 8, 29; Ch 17.1, 17.3

Wherever HI dominates the ISM, all atoms are found in the **ground state** (^{2}S)(n=1). The next excited level (^{2}P) is about 10 eV above the ground-level. This excitation is very rare and and quickly falls back to the ground level, so there is no significant population of this level. For example, consider potential excitation mechanism: collision of H-atom with cosmic ray particle (lots of energy, probability is once per 10^{17} seconds), ionizes the H-atom. Recombination results in some atoms winding up in ^{2}P state. But A coefficient for spontaneous emission from 2P $\rightarrow$ 2S is 10^8 s^{-1}. Hence, this excitation process results in ralative population of 2P level of $10^{-8}/10^{17} = 10^{-25}$!

ISM is too cool for collisions to happen often and cosmic rays are rare. Possible tracers of HI gas:

1. 21 cm HI transition (=hyperfine transition) in emission or absorption.

2. Lyman absorption lines against hot background stars.

Only the ^{2}S level is populated. HI is hard to find in the ground state; fine structure $\rightarrow$ different angular momentum.

See Draine Ch 4 (& 5) on notation of energy levels and atomic structure.
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Excitation and radiative transport for the 21-cm line

Spin of proton and electron:

- Parallel (upper energy)
- Anti-parallel (lower energy)

(Spin is around particle's own axis, not to be confused with angular momentum). Motions specified by maxwellian velocity distribution, and collisions dominate the level populations (excite and de-excite).
Energy difference (very small):

$$h\nu = 9.4 \times 10^{-18}\,\mathrm{erg}$$
$$\nu = 1420.4\,\mathrm{MHz}$$
$$\lambda = 21.11\,\mathrm{cm}$$

Spontaneous emission probability is *very* small:

$$A_{kj} = 2.86 \times 10^{-15}\,\mathrm{sec}^{-1}$$

$$\rightarrow \text{lifetime} = 1.10 \times 10^7 \text{ years}$$

More frequently, atoms will flip energy states by *collisions*. These dominate transitions and cause energy levels to flip.
$\mathrm{H} + \mathrm{H} \rightarrow \mathrm{H}_2^* \rightarrow \mathrm{H} + \mathrm{H}$
excited H_2 molecule (not stable). Chance for collisional excitation per second: $\gamma n(H)$ [s^{-1}].

$\gamma[\times 10^{-11}\ \mathrm{cm^3 s^{-1}}]$	T [K]
0.23	10
3.0	30
9.5	100
16	300
25	1000

As long as $n(H^o) > 10^{-2}\ \mathrm{cm}^{-3}$, $\gamma n_H >> A$, so collisions determine excitation and de-excitation...implies detailed balancing, so Boltzmann distribution is valid for level populations.

$$\frac{n_1}{n_o} = \frac{g_1}{g_o}\exp(-h\nu/kT_k) \approx \frac{g_1}{g_o}(1 - \frac{h\nu}{kT_k}) \approx \frac{g_1}{g_o}$$

where T_k = kinetic gas temperature.

-35- Derivation: ...In general, in situations where stimulated emission and absorption can be neglected, we would obtain

$$\frac{b_1}{b_o} = \frac{1}{1 + \frac{A_{10}}{n_H \gamma_{10}}} \approx 1$$

Note: Potential other excitation process of upper level for HI: HI + Lyα photon $\rightarrow$ n=2 level (2P). 2P level might cascade down to upper 2S hyperfine level.

D 17.3 In principle, one might selectively populate the upper level with this so-called photon pumping. It turns out that, due to thermilization of the photons this anomalous population of the levels does not generally occur. (Lyα photons frequently scatter off A-atom again = LTE situation).

-37- Since the b's are effectively 1, our main result is that

Are the HI levels always populated mainly by collisions? No, if n(H^o drops low enough, and if the HI is warm it doesn't work as well.

The warm HI in our Galaxy has $n_H \sim 0.4$ cm $^{-3}$ and $T_k \approx 8000$ K. But T_s = spin temperature, of order $T_k/5$, if it weren't for excitation by Lyα photons.

But it is true that the 3K background radiation field does *not* significantly disturb the equilibrium set up by collisions, as we discussed above.

If n_H drops to very low values, collisional excitation is ineffective. But even in that case, $T_s \approx T_k$ because of the above mentioned Lyα excitation.

In any case, in general we now obtain for the emission coefficient:

which is independent of T in most circumstances, and for the absorption coefficient:

Simple case of a single layer of gas

IX. Atomic Structure

- electron spin -I6-
- spin-orbit coupling -I8-
- atoms with multiple electrons -I10-
- transition rules
- x-ray emission, Zeeman effect -I20-

X. HII Regions -51-

- Stromgren Theory -52-
- HII Region spectra
  - Continuum Radiation -63-

XIV. Heating and Cooling -151-

- General
  - Primary heat source: photoionization -152-
  - Primary cooling source: inelastic collisions -153-
- HII regions (ionized) -154-
  - Heating -154-
  - Cooling -155-
- HI gas (neutral) -160-
  - Dominant cooling line from CII (IP = 11.26 eV)
  - HI naturally cool, but observe very warm HI!
  - General players:
    * Cooling function -160-
    * Heating function -161-