Physics 224
The Interstellar Medium

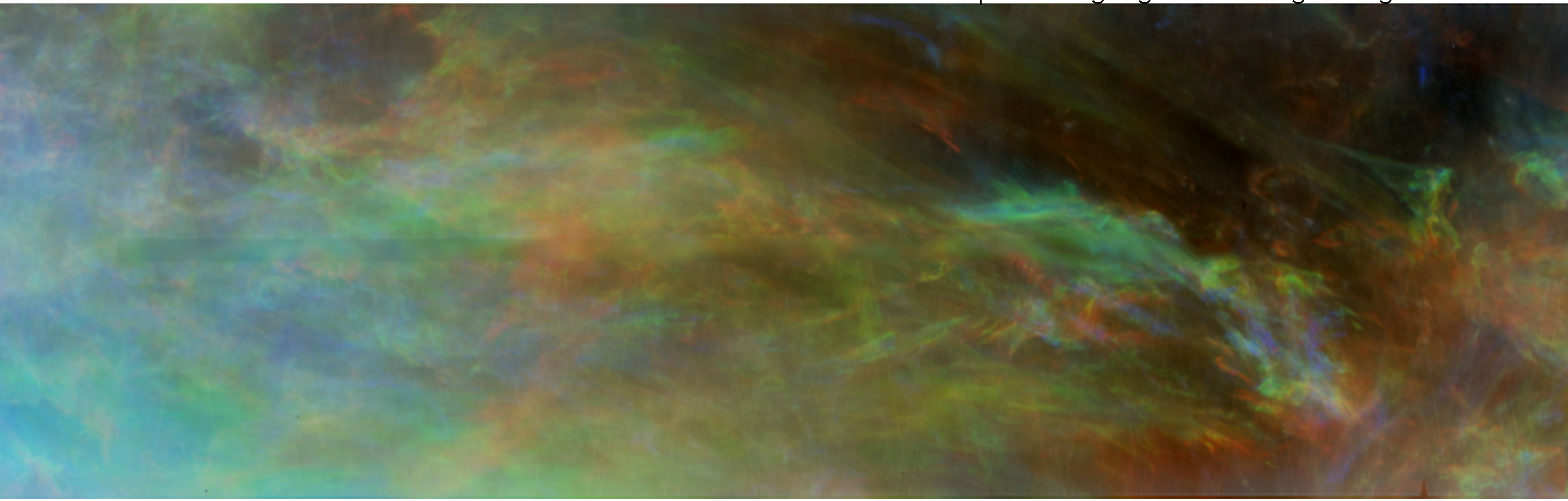
Lecture #13: Neutral Gas, Photodissociation Regions & the HI to H_2 transition

Outline

- Part I: Neutral Gas
- Part II: HI to H_2 Transition
- Part III: Photodissociation Regions

Is the FGH model a good representation of the ISM?

https://sites.google.com/site/galfahi/galfa-hi-science



part of the GALFA HI Survey
colors = different velocity ranges

Is the FGH model a good representation of the ISM?

Do we expect to find much gas in the unstable region?

Compare thermal and dynamical timescales:

$$\tau_{\rm cool} = \frac{nkT}{\Lambda}$$

nkT ⟵ thermal energy density = pressure

Λ ⟵ cooling rate per unit volume

* note same for heating since $\Gamma = \Lambda$

$\tau_{\rm cool}$ ~ 0.1 Myr for unstable gas with T ~ 2000 K and n ~ 1.5 cm^{-3}

Is the FGH model a good representation of the ISM?

Do we expect to find much gas in the unstable region?

Compare thermal and dynamical timescales:

$$\tau_{\rm dyn} \sim \frac{L}{c_s}$$

where sound speed: $$c_s = \sqrt{\frac{kT}{m}}$$

$$\tau_{\rm dyn} \sim 6.7{\rm Myr}\left(\frac{L}{1{\rm pc}}\right) T^{-1/2}$$

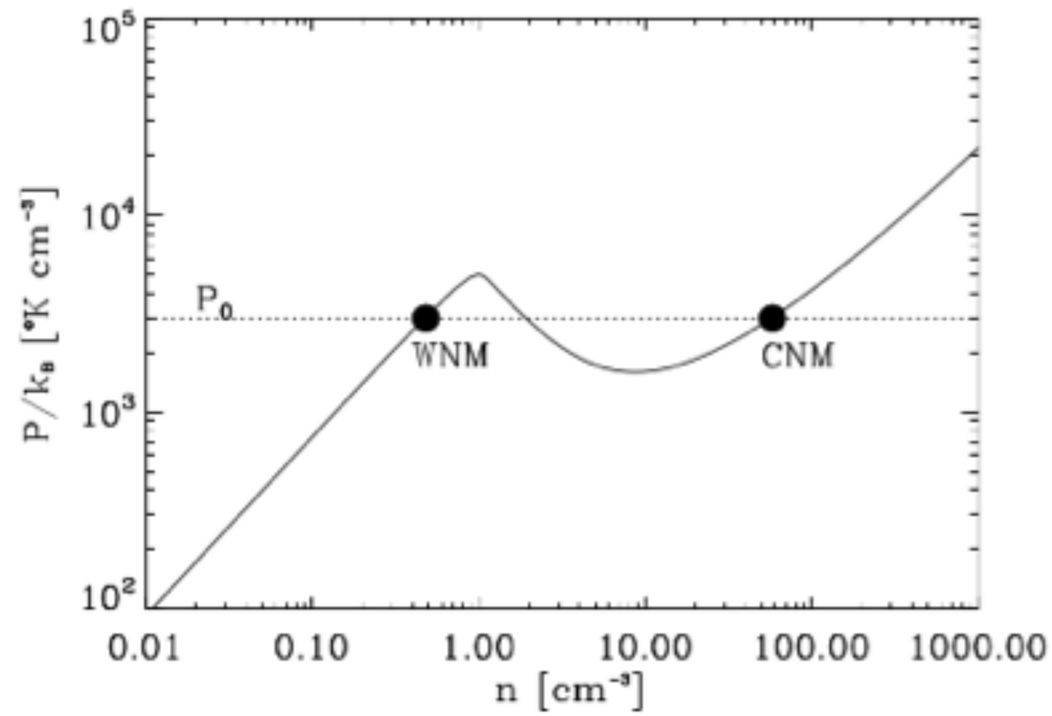
For L~10 pc, T~2000 K

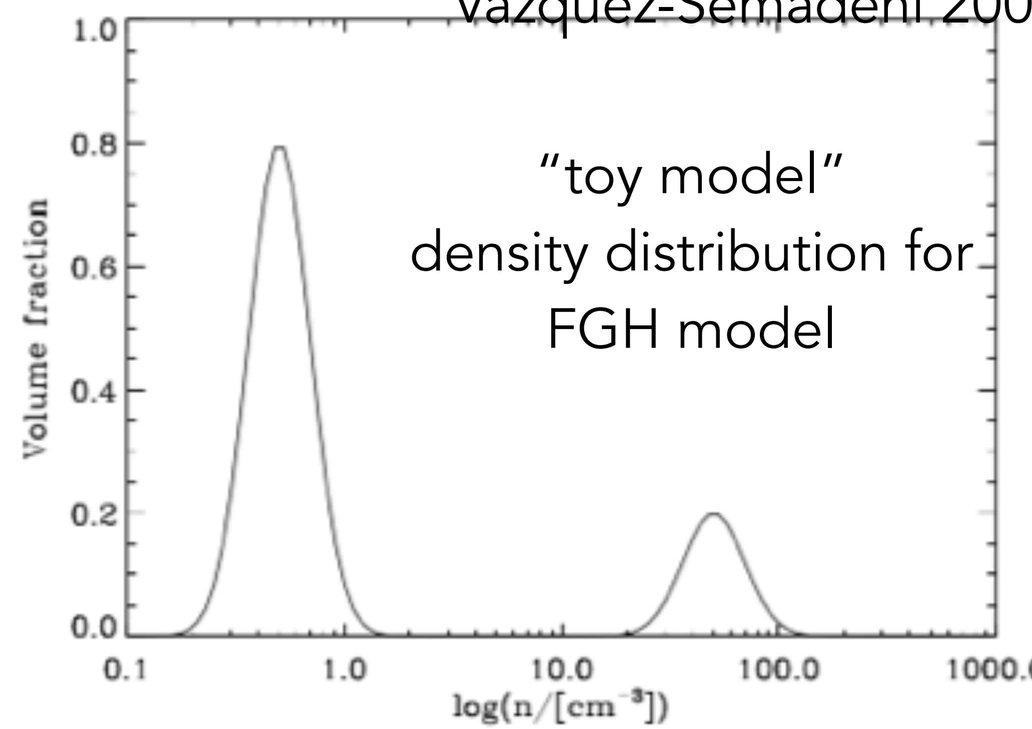
$\tau_{\rm dyn}$ ~ 1.5 Myr

Unstable gas should cool quickly relative to dynamical time.

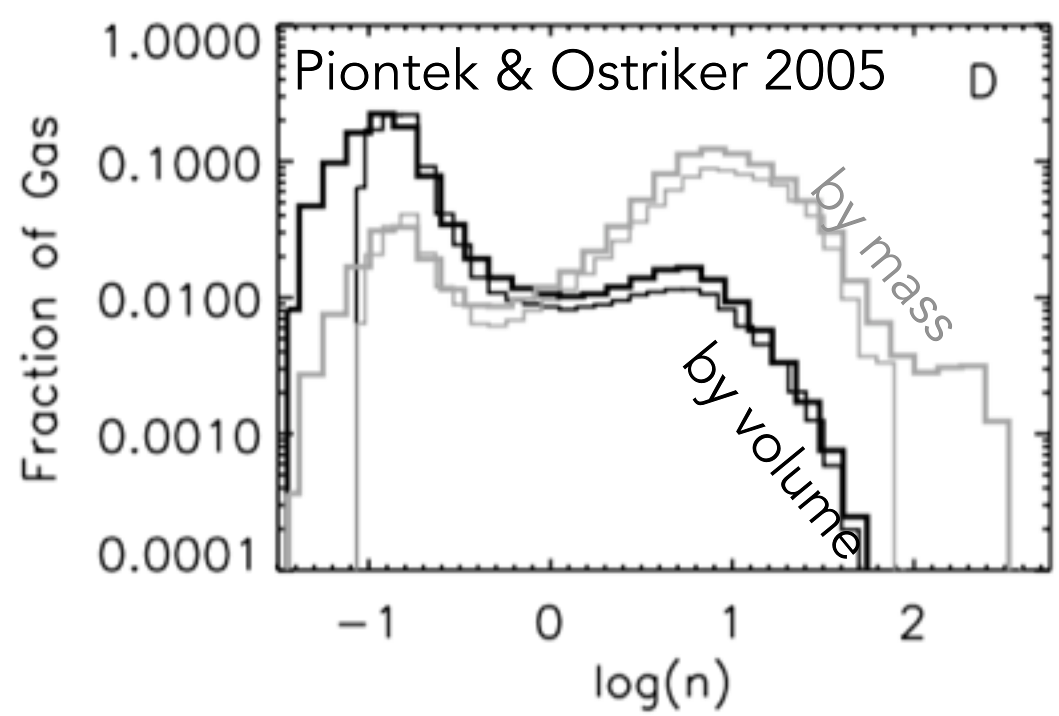
Is the FGH model a good representation of the ISM?

Vazquez-Semadeni 2009

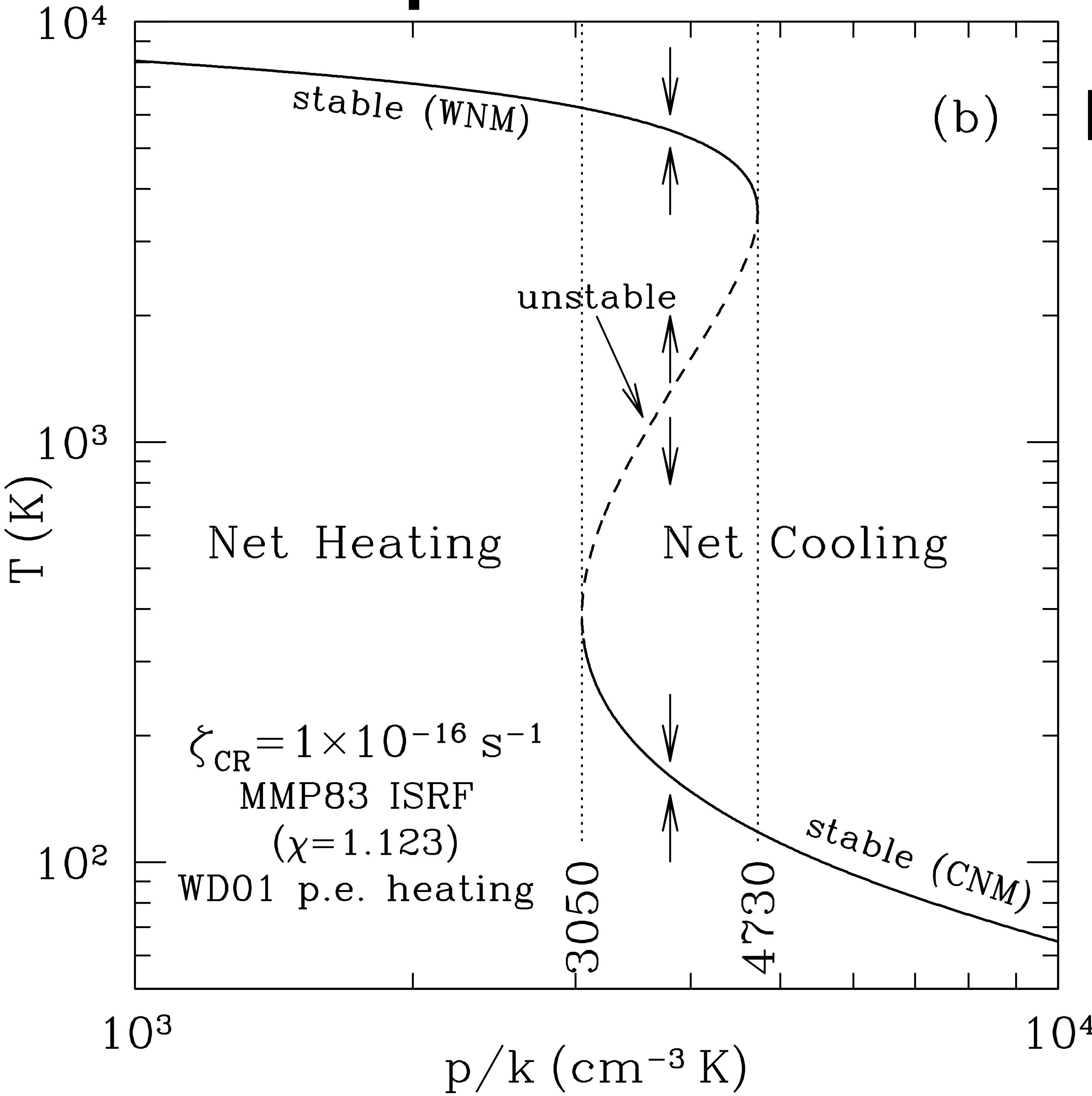




Simulations with turbulence suggest substantial amounts of gas between F&H phases



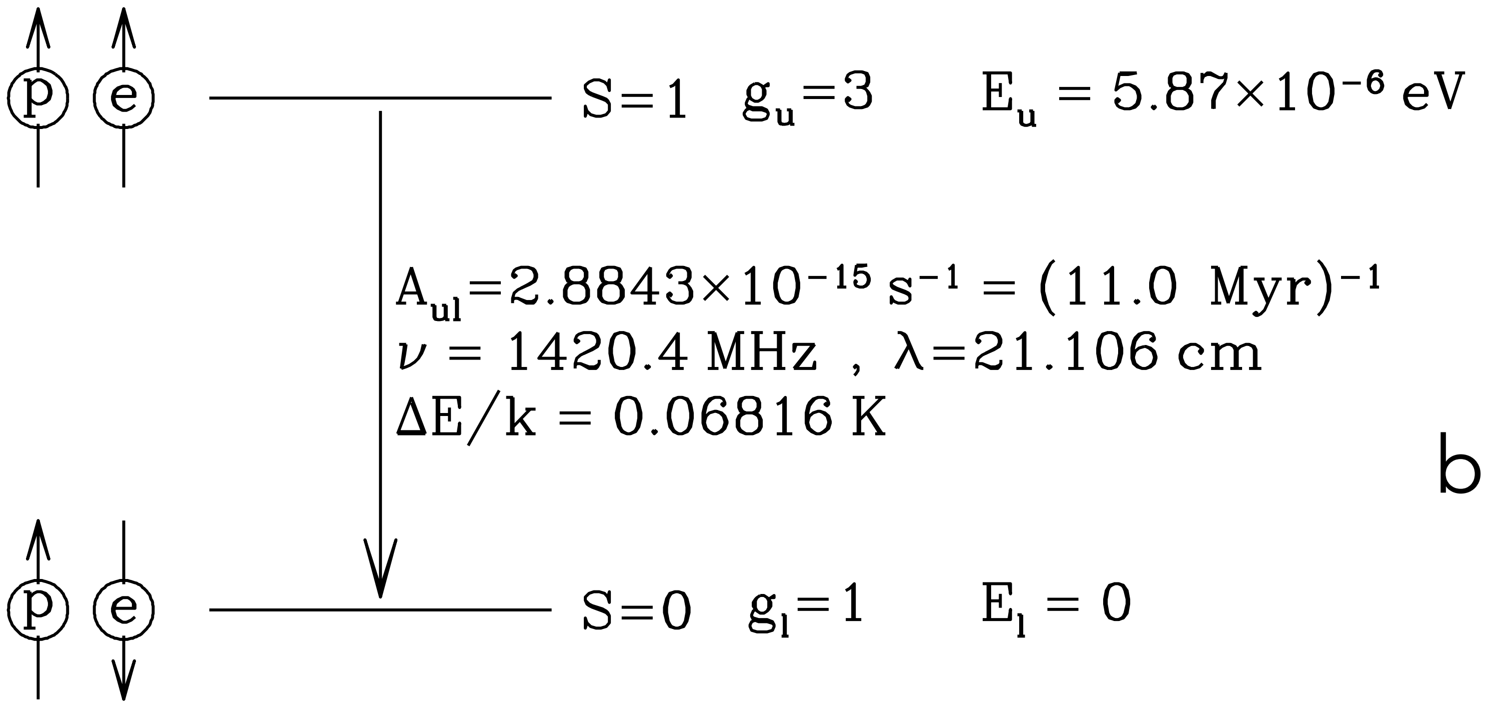
Is the FGH model a good representation of the ISM?



How can we test this model?

Measure the n & T of HI gas and see if it matches the predicted n,T ranges for CNM and WNM stable phases.

HI Spin Temperature

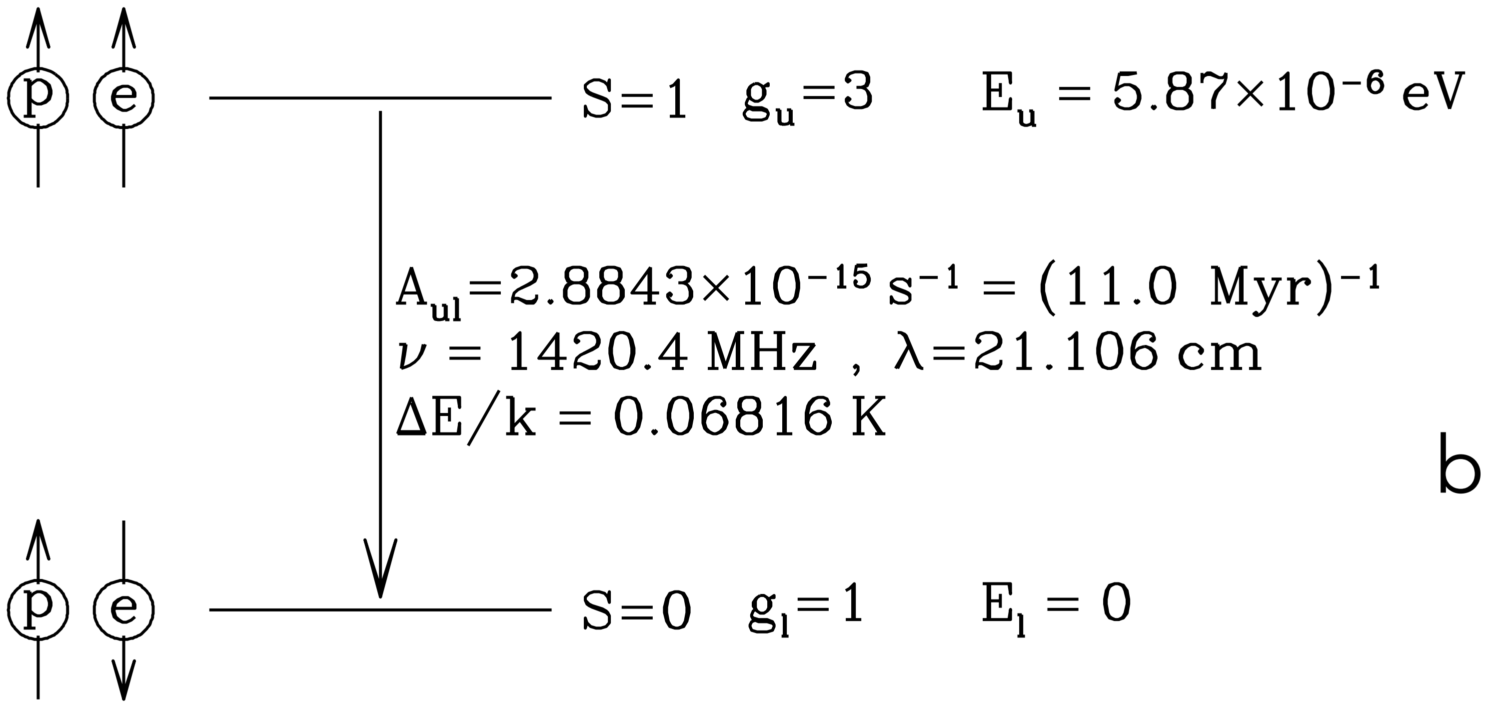


$T_{exc} \equiv T_{spin} \gg 0.0682$ K because cosmic microwave background can populate levels

Under most ISM conditions, 75% of HI is in upper level. *Emissivity is independent of T_{spin}!!*

$$j_\nu = n_u \frac{A_{ul}}{4\pi} h\nu_{ul}\phi_\nu = \frac{3}{16\pi} A_{ul}\ h\nu_{ul}\ n(\mathrm{H\ I})\ \phi_\nu$$

HI Spin Temperature



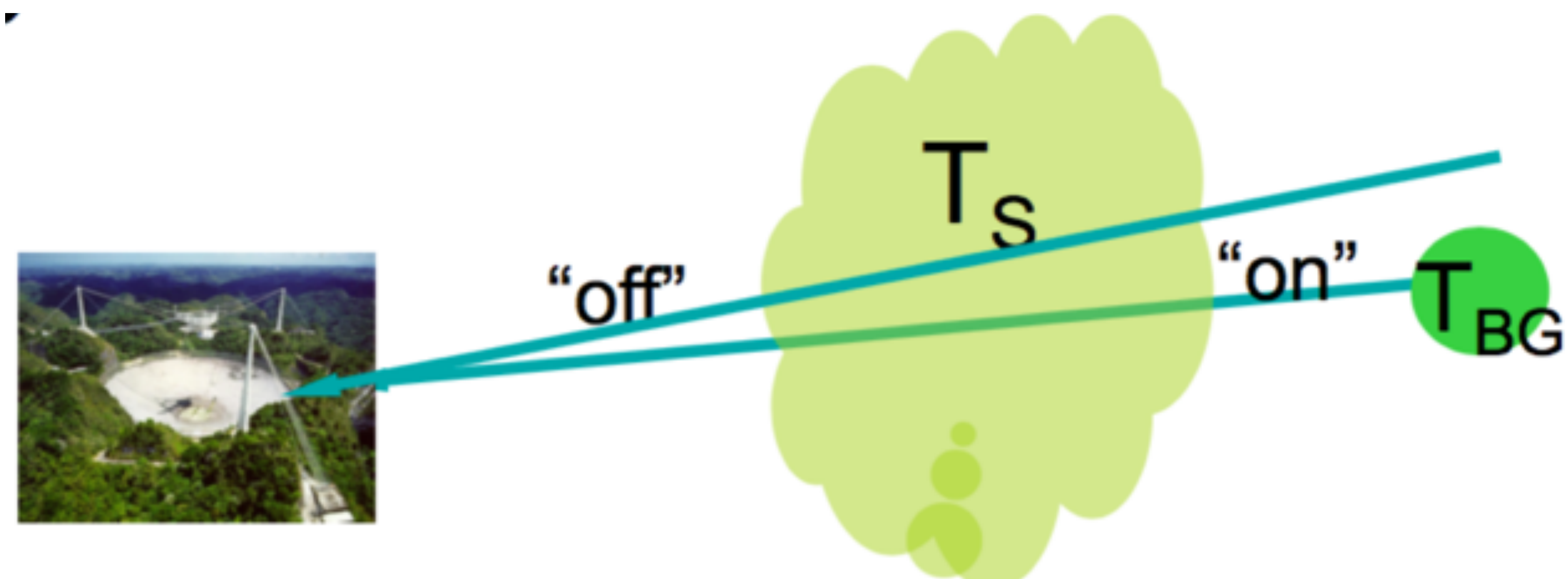
$T_{exc} \equiv T_{spin} \gg 0.0682$ K
because cosmic microwave background can populate levels

absorption coefficient depends inversely on T_{spin}
as a consequence of stimulated emission not being negligible!

$$\kappa_\nu \approx \frac{3}{32\pi} A_{ul} \frac{hc\lambda_{ul}}{kT_{spin}} n(\mathrm{H\ I})\phi_\nu$$

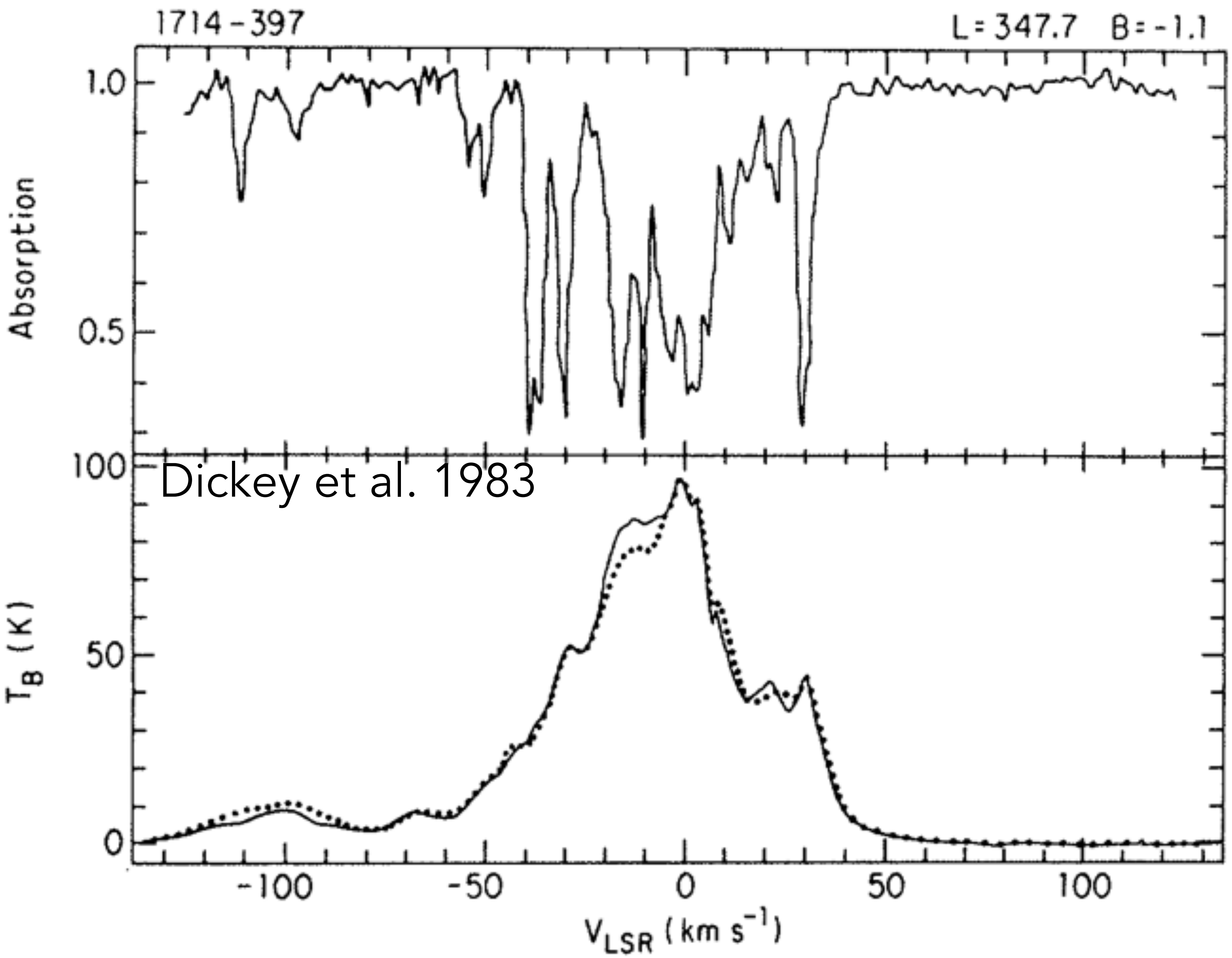
HI Spin Temperature

Measuring spin temperature



$$T_b^{on} = T_{bg} e^{-\tau} + T_s(1 - e^{-\tau})$$
$$T_b^{off} = T_s(1 - e^{-\tau}) \quad (1)$$

HI Spin Temperature

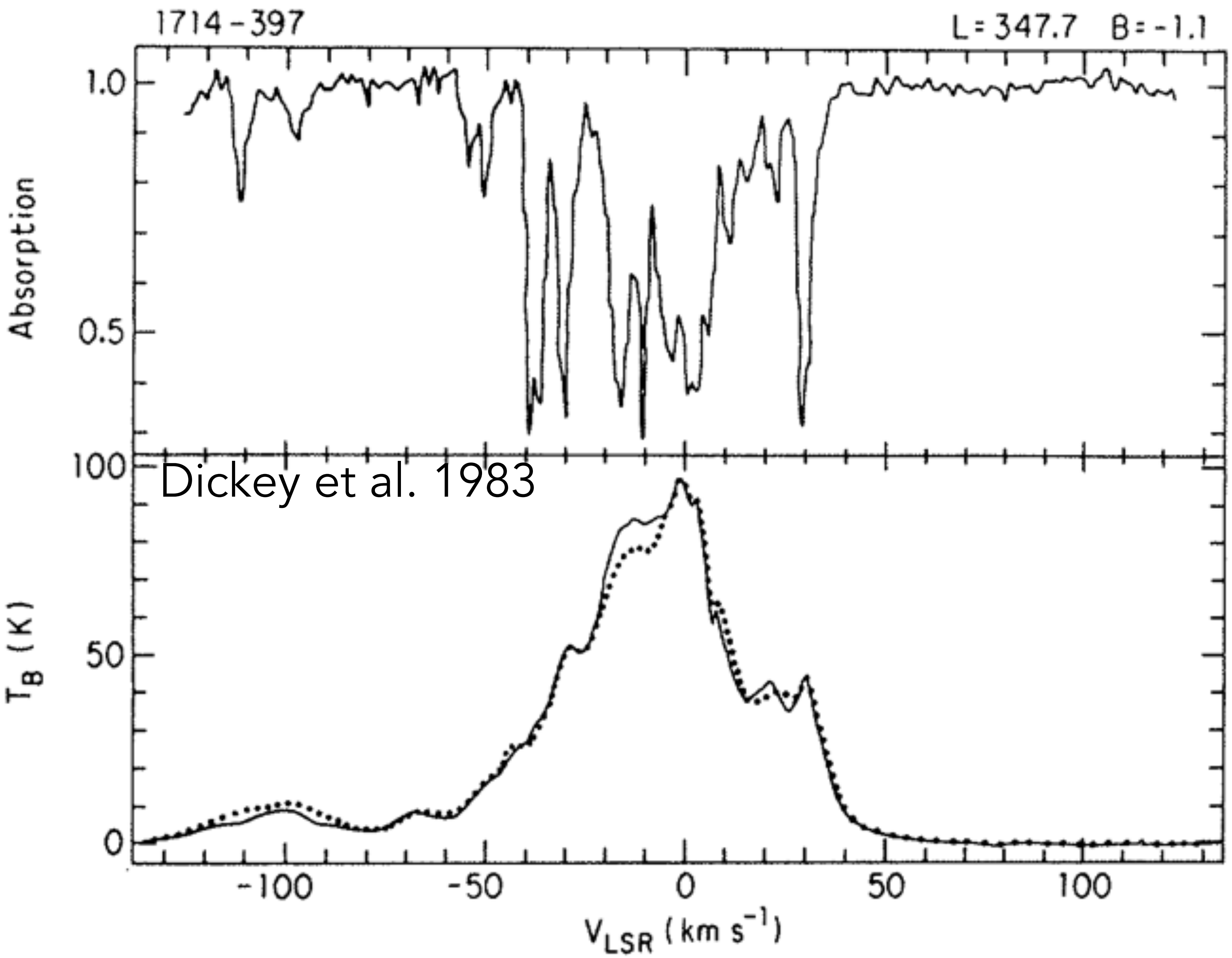


Absorption - weighted to low T

Emission - independent of T

$\langle T_{spin} \rangle = T_B/(1-e^{-\tau})$

HI Spin Temperature

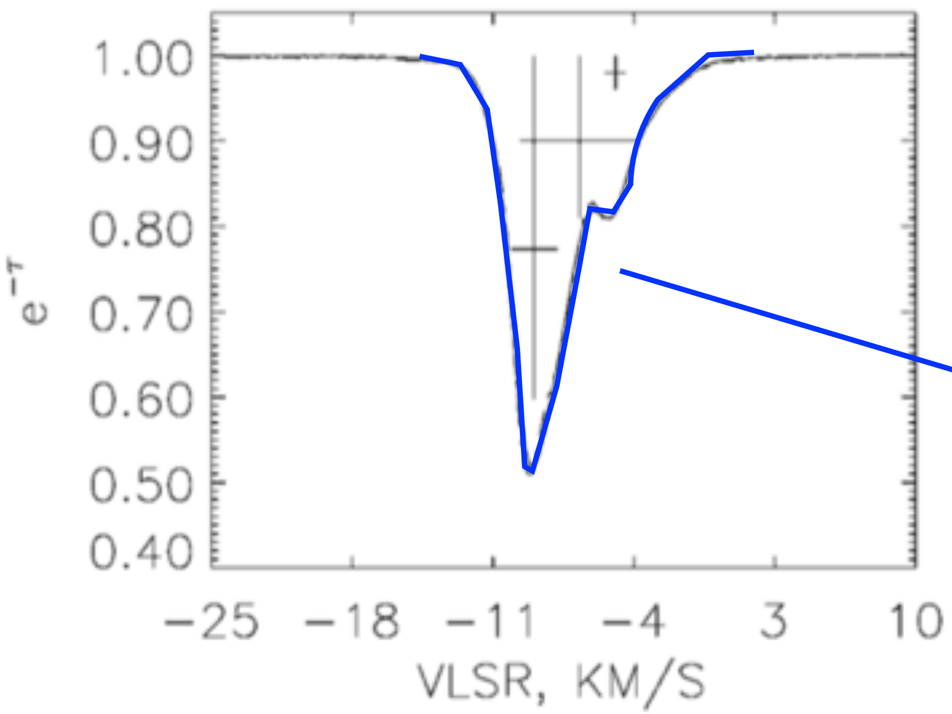


Assume T_{WNM} is too big to contribute much to the absorption.

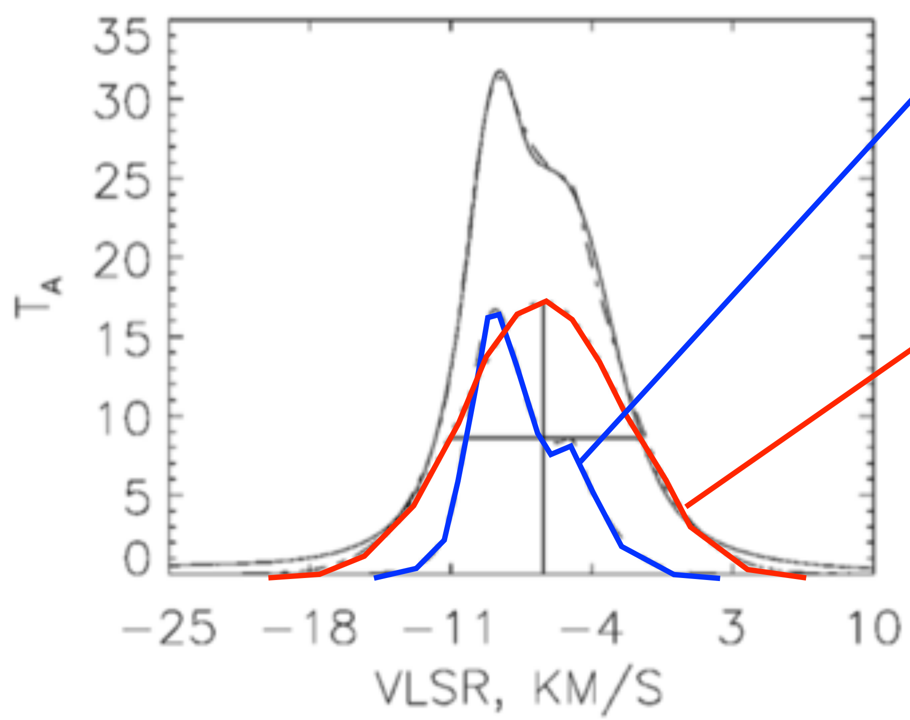
$\tau \sim n_{CNM}/T_{CNM}$

$T_B \sim n_{CNM} + n_{WNM}$

Observed HI Spin Temperature



Heiles 2001



Assume CNM dominates absorption.

Fit absorption component and emission component with same Gaussian components (σ_V) to get N_{CNM}, T_{CNM}

Fit emission component with additional Gaussian and N_{WNM}.

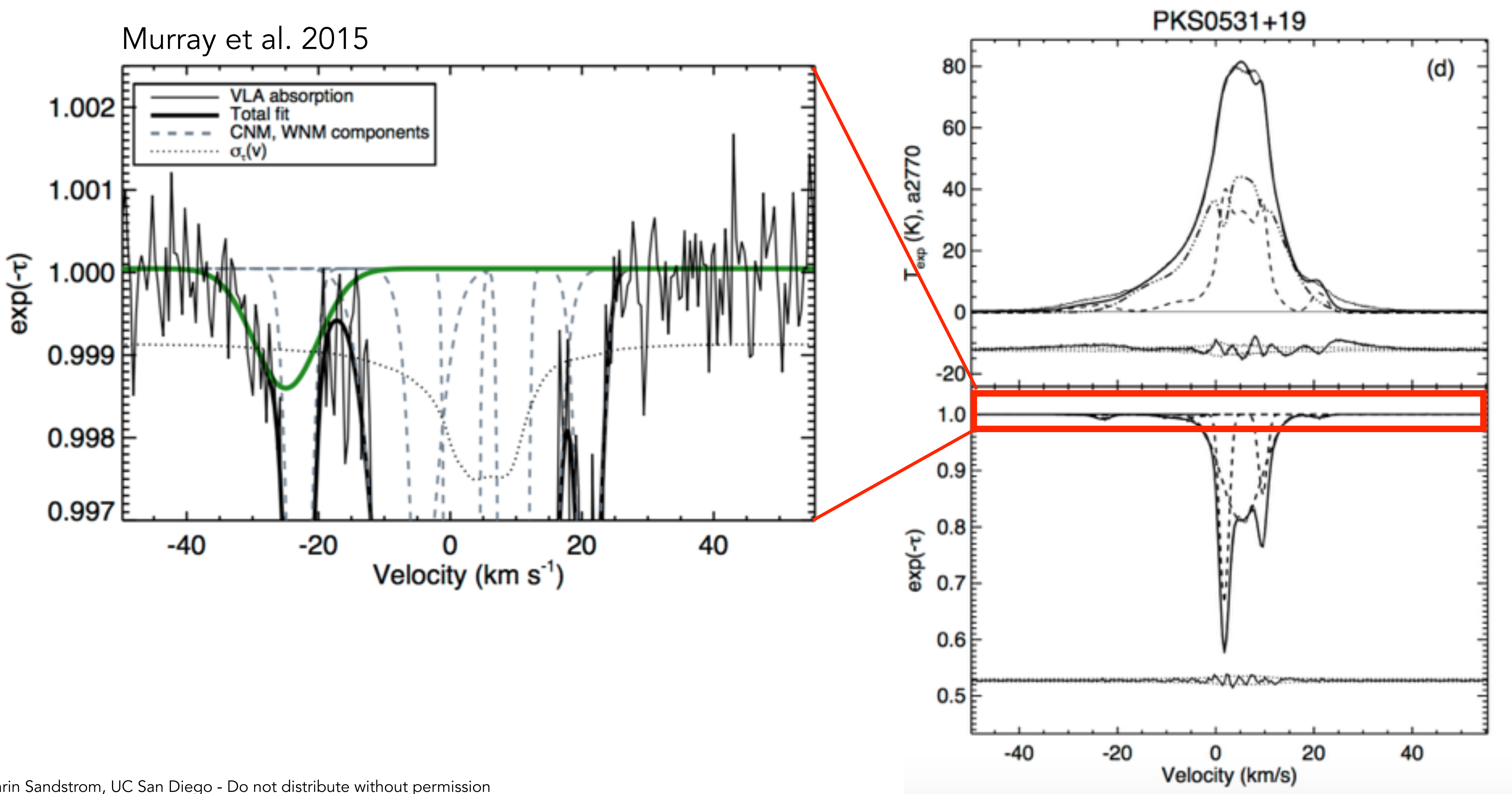
Get upper limit on T_{WNM} from velocity width (upper limit because of turbulent contribution).

Get lower limit on T_{WNM} from residual absorption.

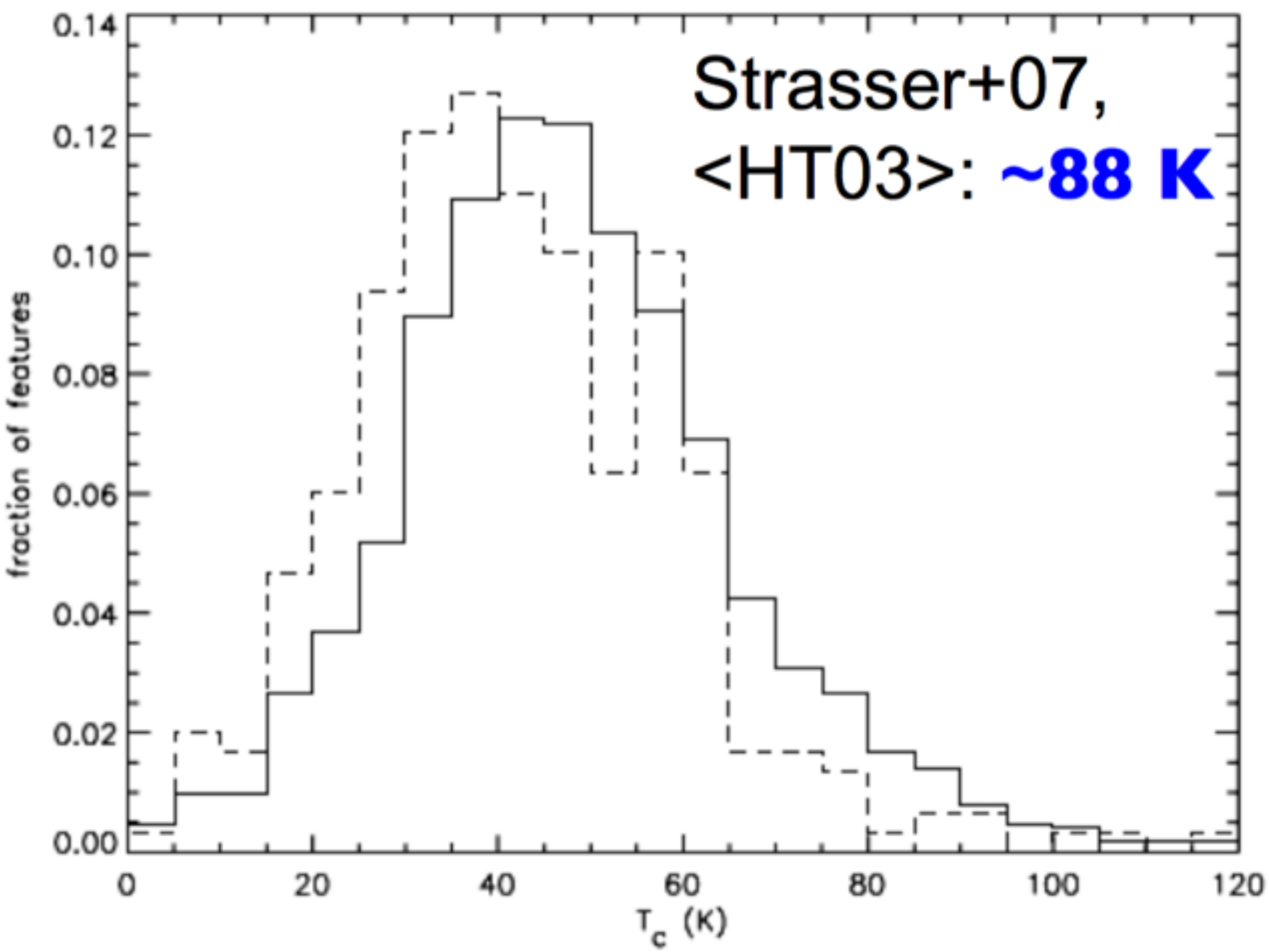
Observed HI Spin Temperature

Measuring absorption from the WNM requires very high S/N measurements.

Murray et al. 2015



Observed HI Spin Temperature

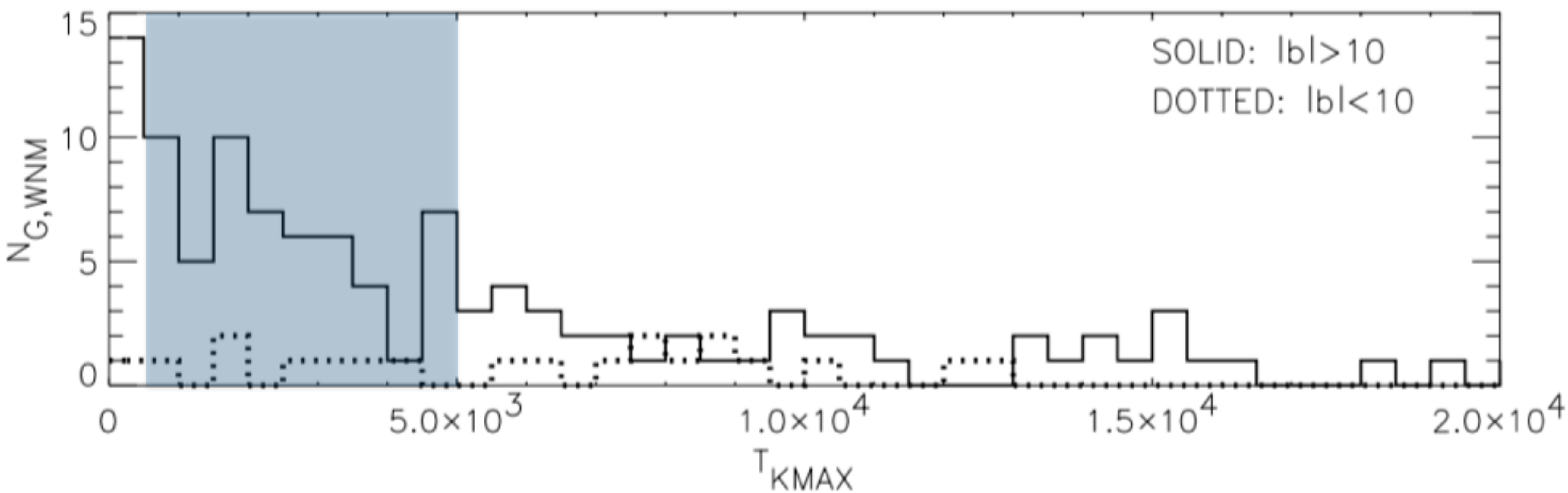


observed CNM components have T ~ 40-80 K

Observed HI Spin Temperature

Evidence for "unstable" phase (500 < T < 5000)

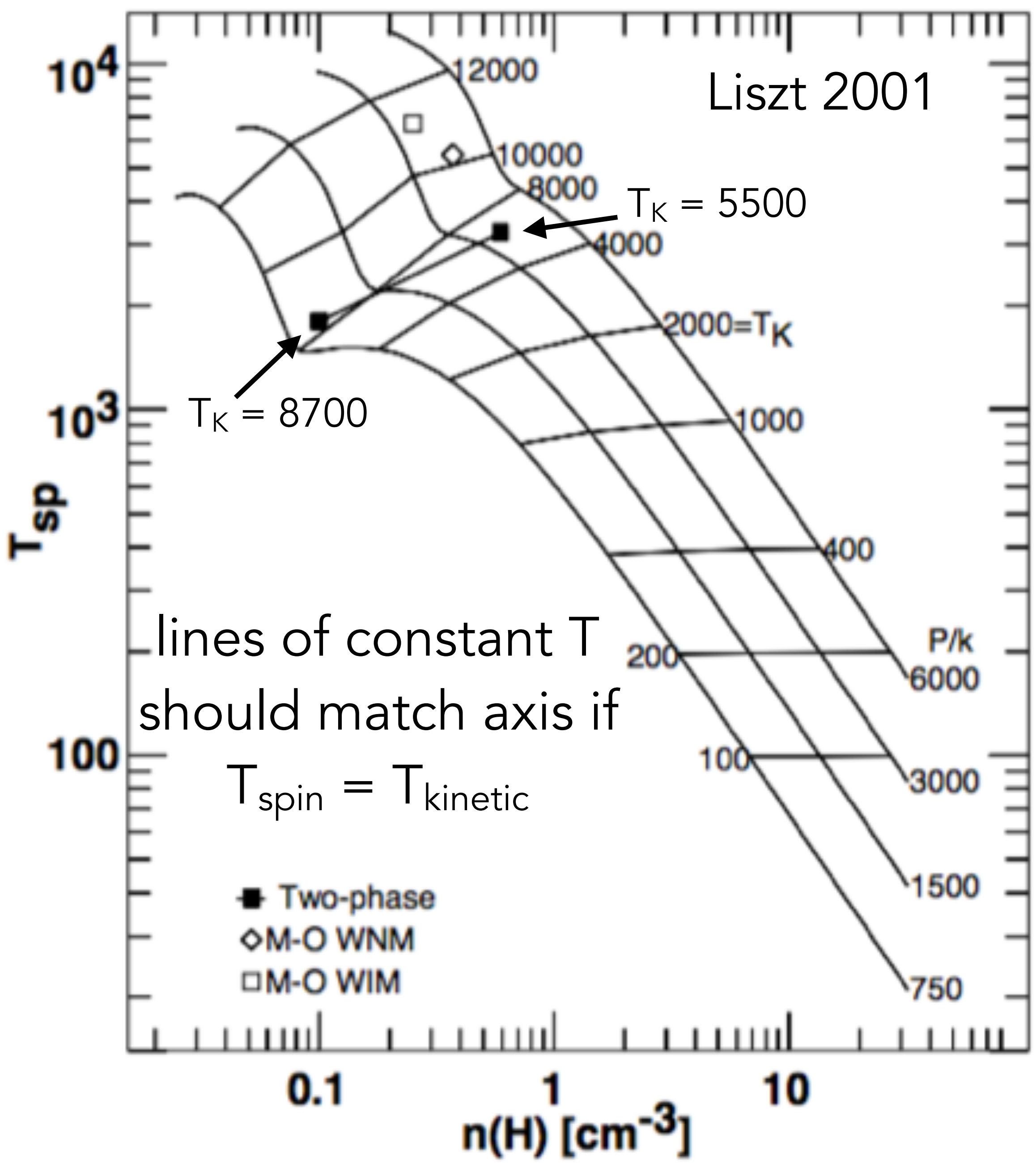
Heiles & Troland 2003



Upper limit on T_{WNM}

Observed HI Spin Temperature

Important wrinkle: thermalization of HI levels in WNM

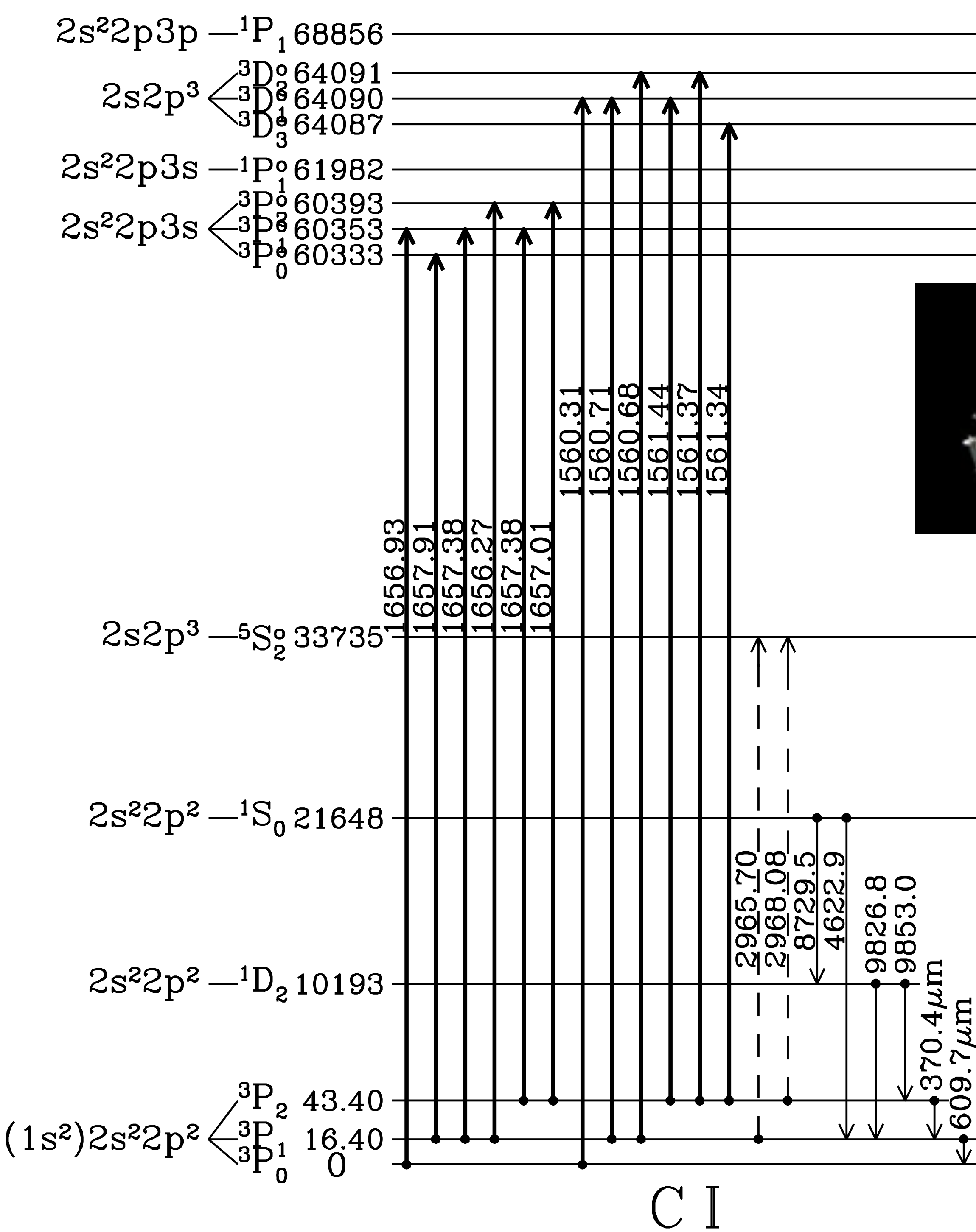


Density in the WNM is too low to thermalize levels to predicted WNM temperatures.

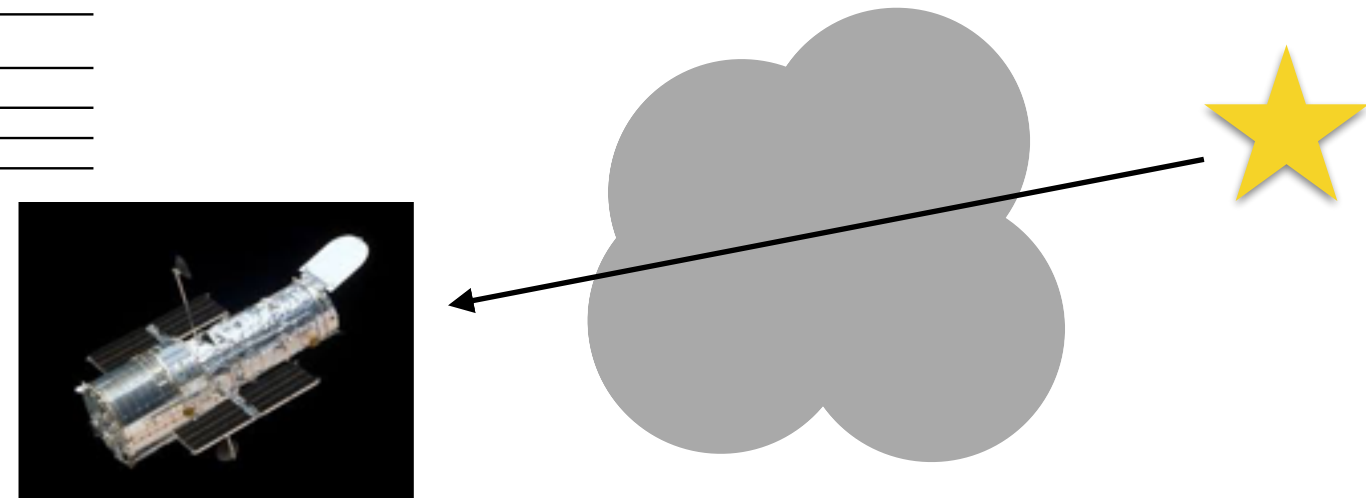
However, scattered Lyα radiation can contribute to thermalizing levels as well.

(Liszt 2001)

Thermal Pressure from [CI]

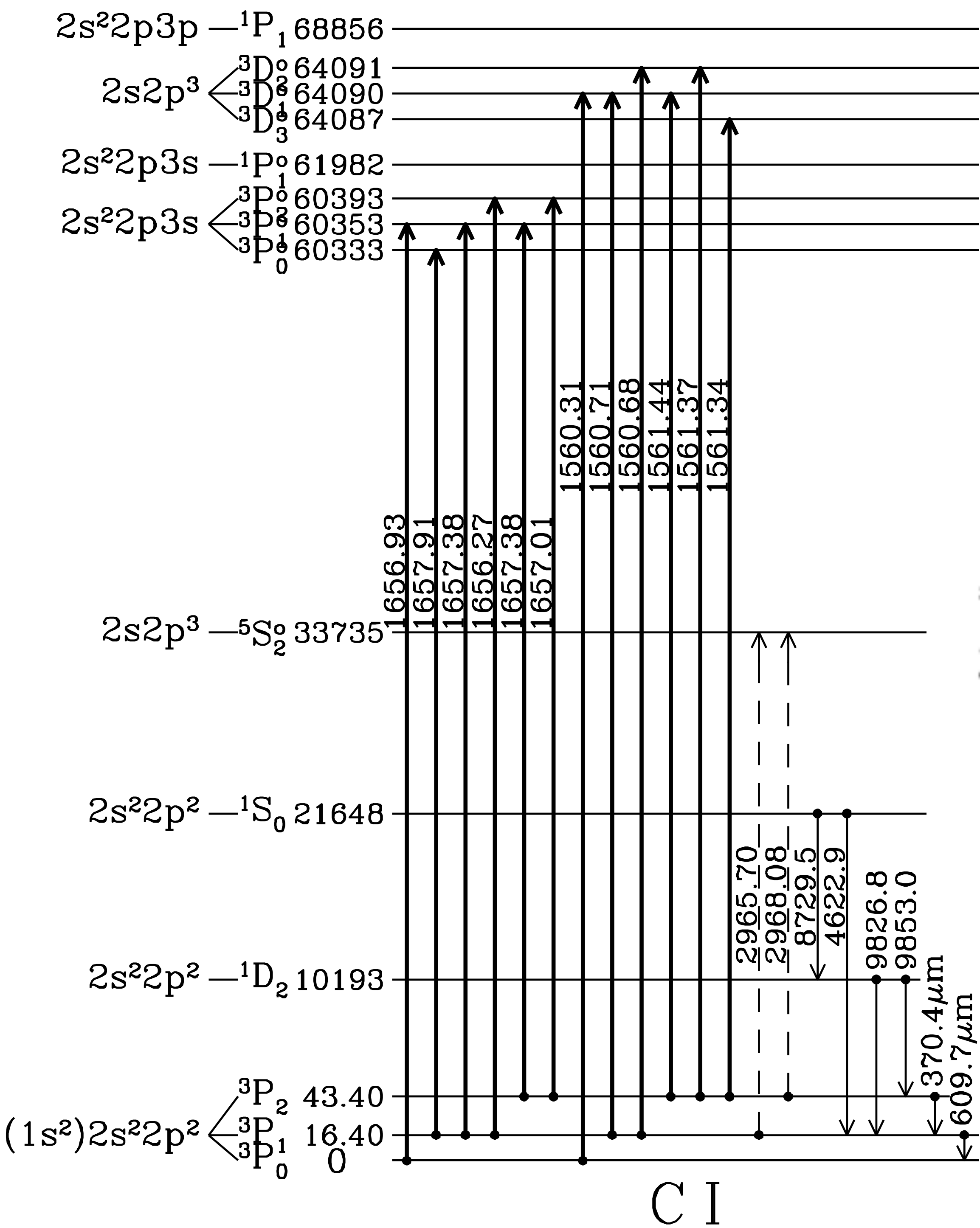


Jenkins & Tripp 2001, 2011



Measure UV absorption lines from ground state, collisionally populated fine structure levels of [CI] - populations set by n,T

Thermal Pressure from [CI]



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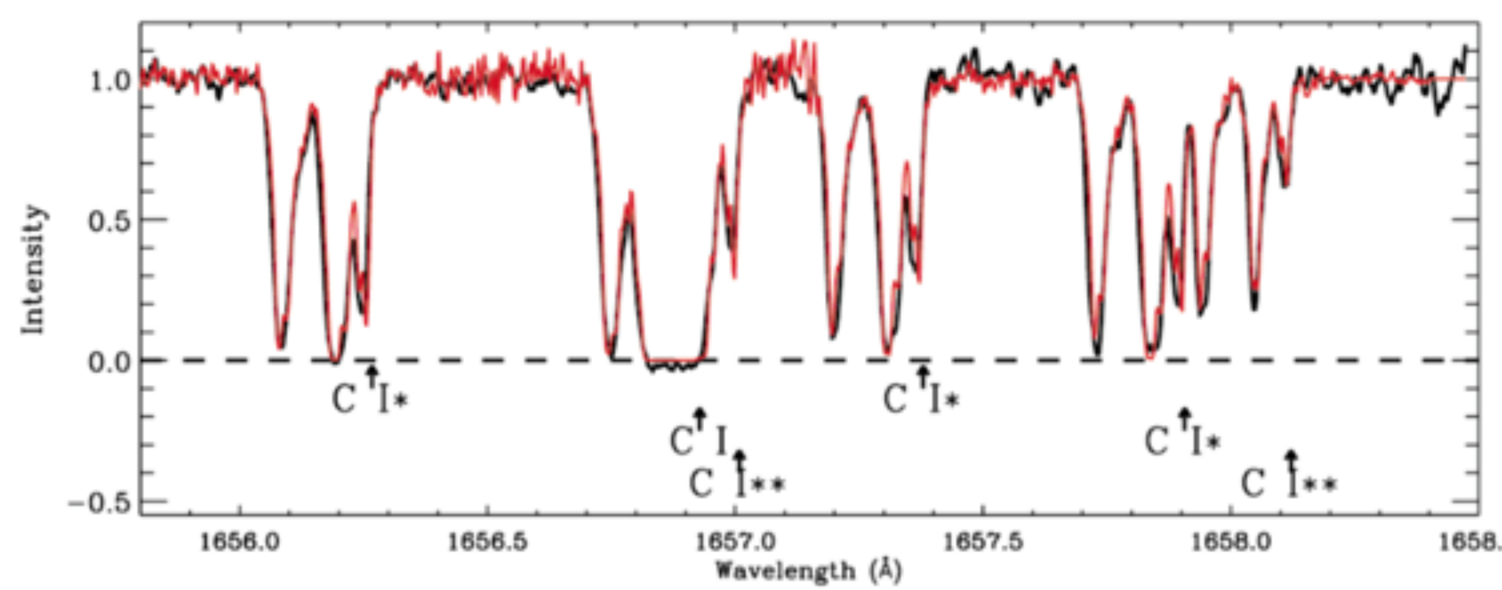

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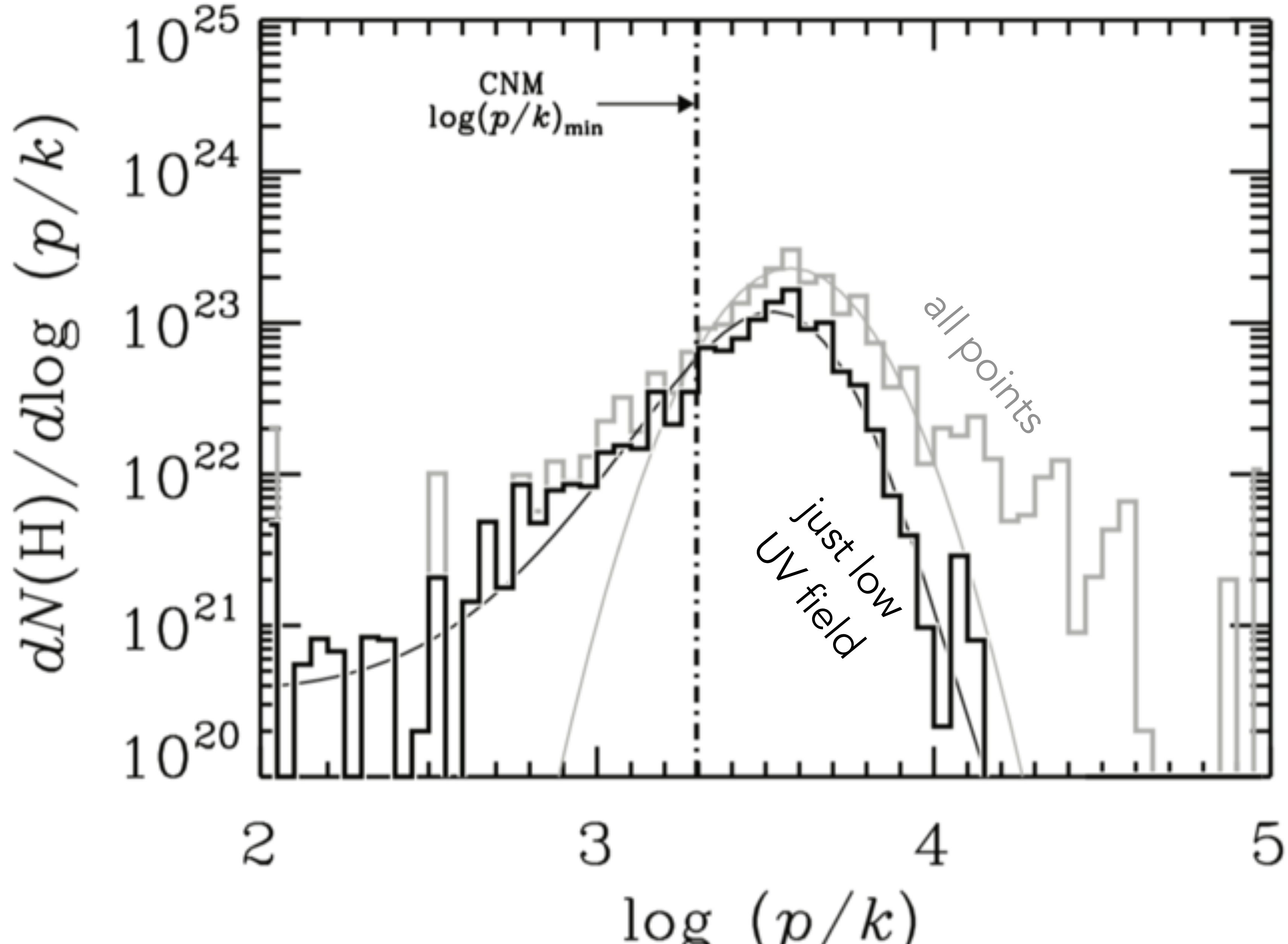
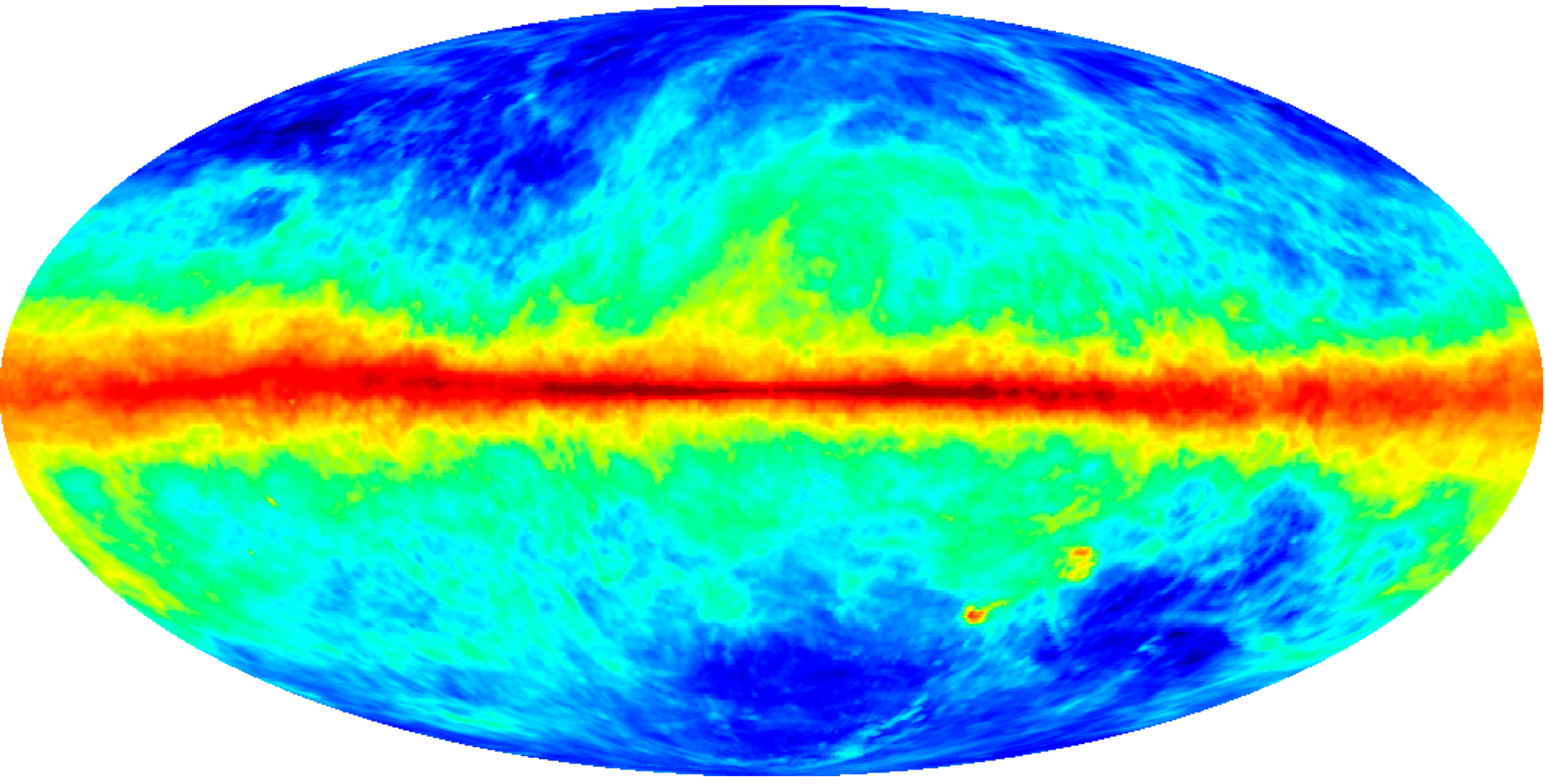


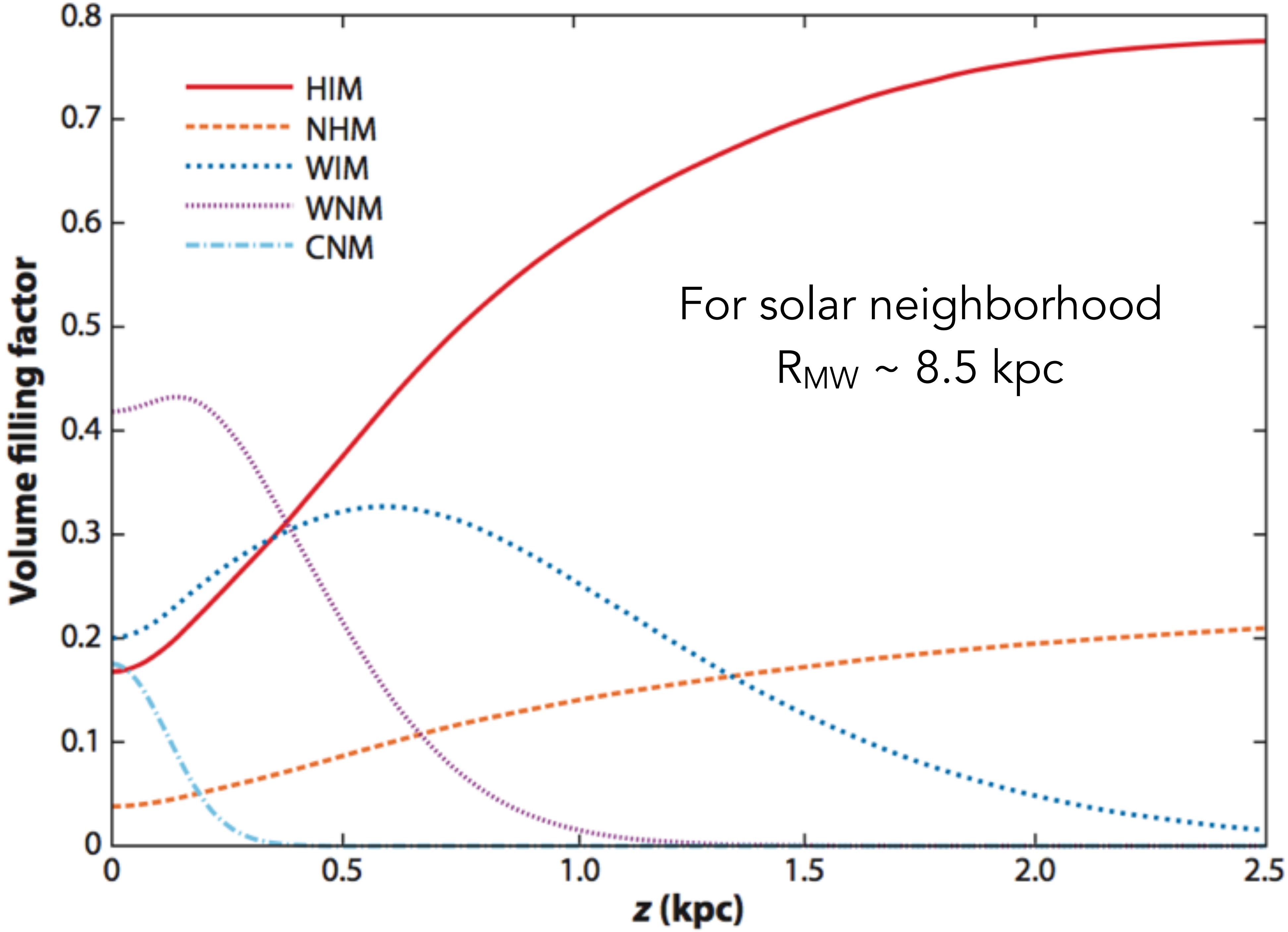
Most gas is at pressures that agree with the FGH picture, but there are tails of low & high pressure that are probably related to turbulence.

All-Sky Map of N(HI) from the Leiden-Argentine-Bonn Survey (Kalberla et al. 2005)



Distribution of HI in the MW

Kalberla & Kerp 2009, ARA&A

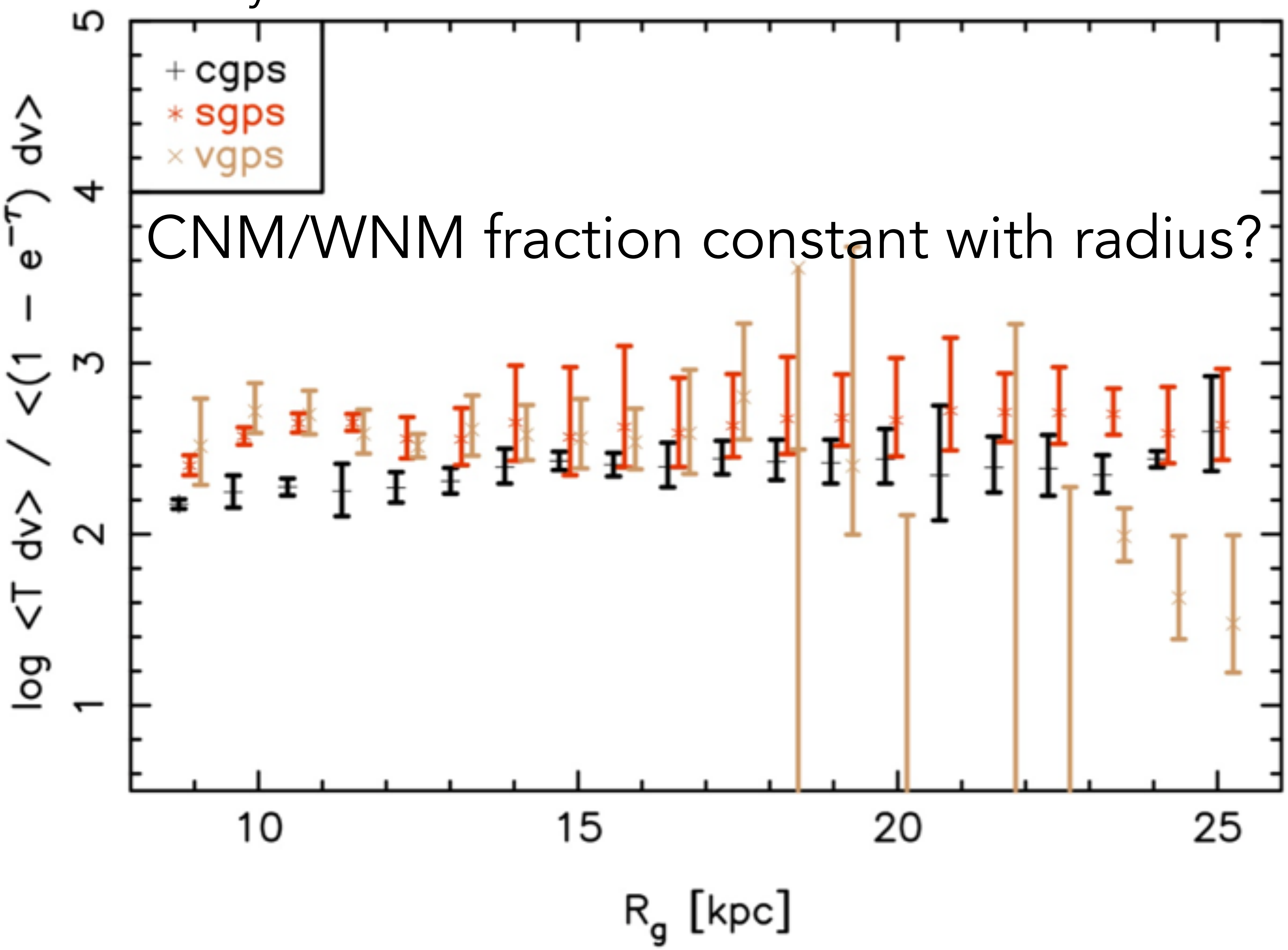


For solar neighborhood
R_{MW} ~ 8.5 kpc

height above Galactic midplane

Distribution of HI in the MW

Dickey et al. 2009



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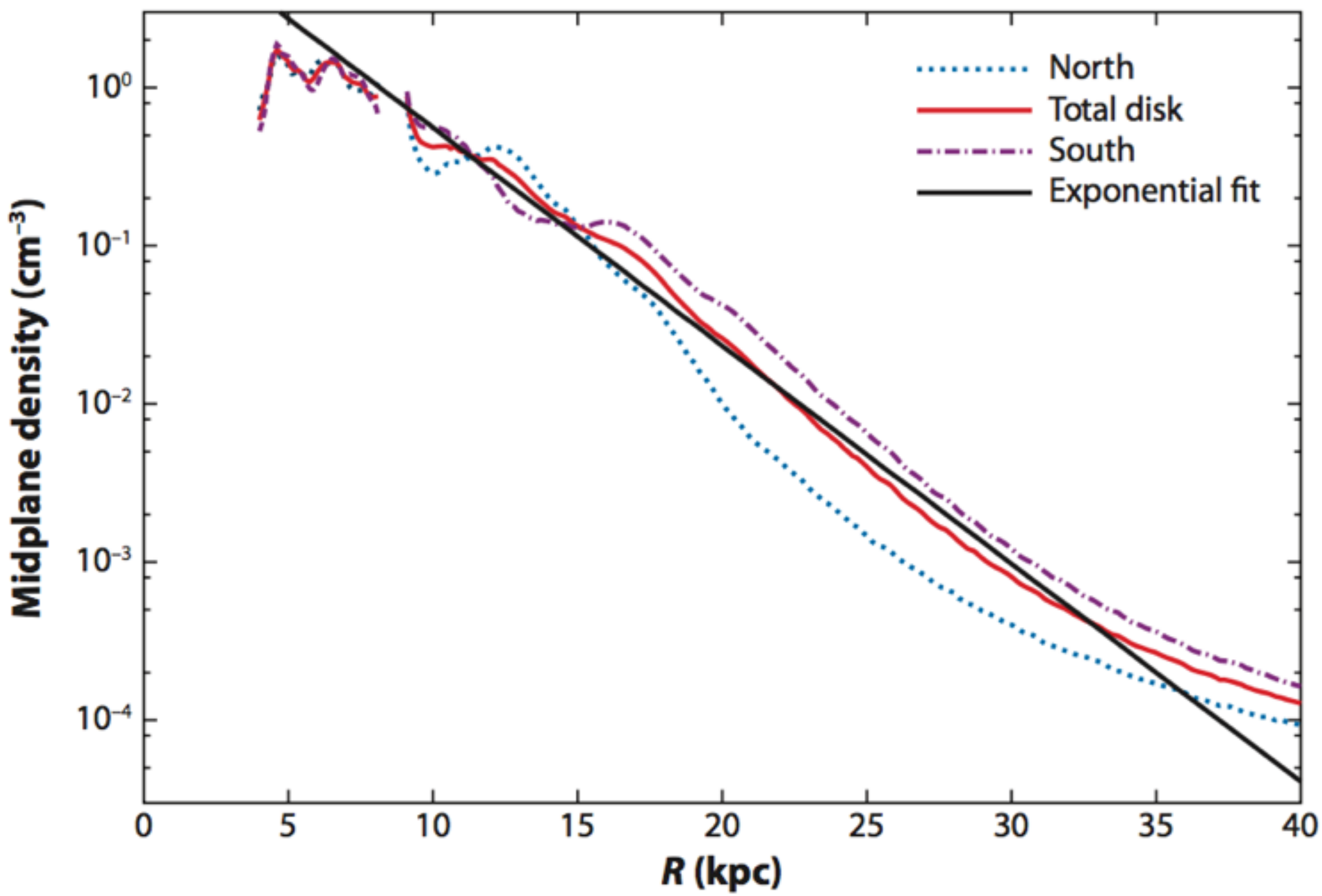
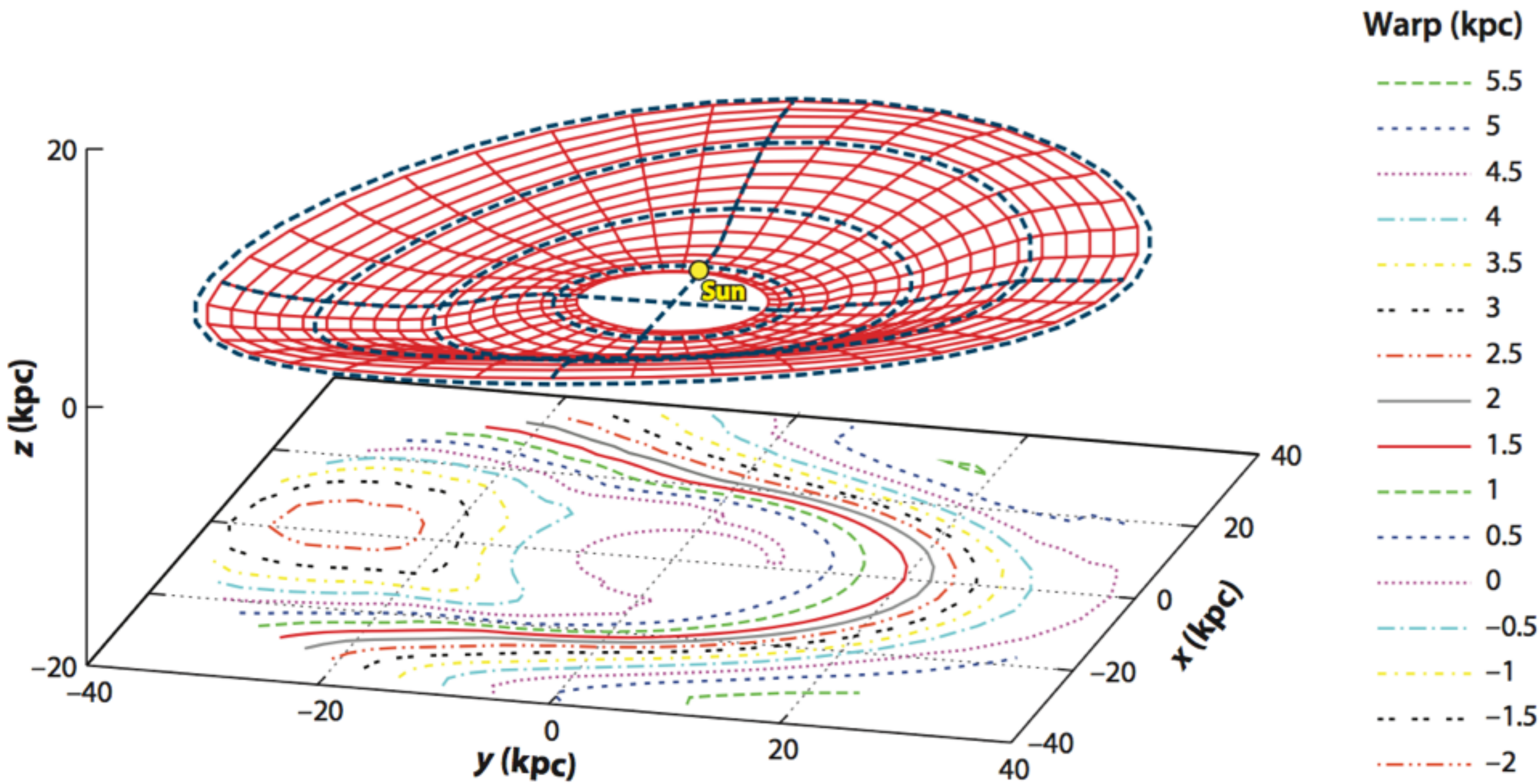
Kalberla & Kerp 2009, ARA&A



distance from Galactic center

Distribution of HI in the MW

Kalberla & Kerp 2009, ARA&A



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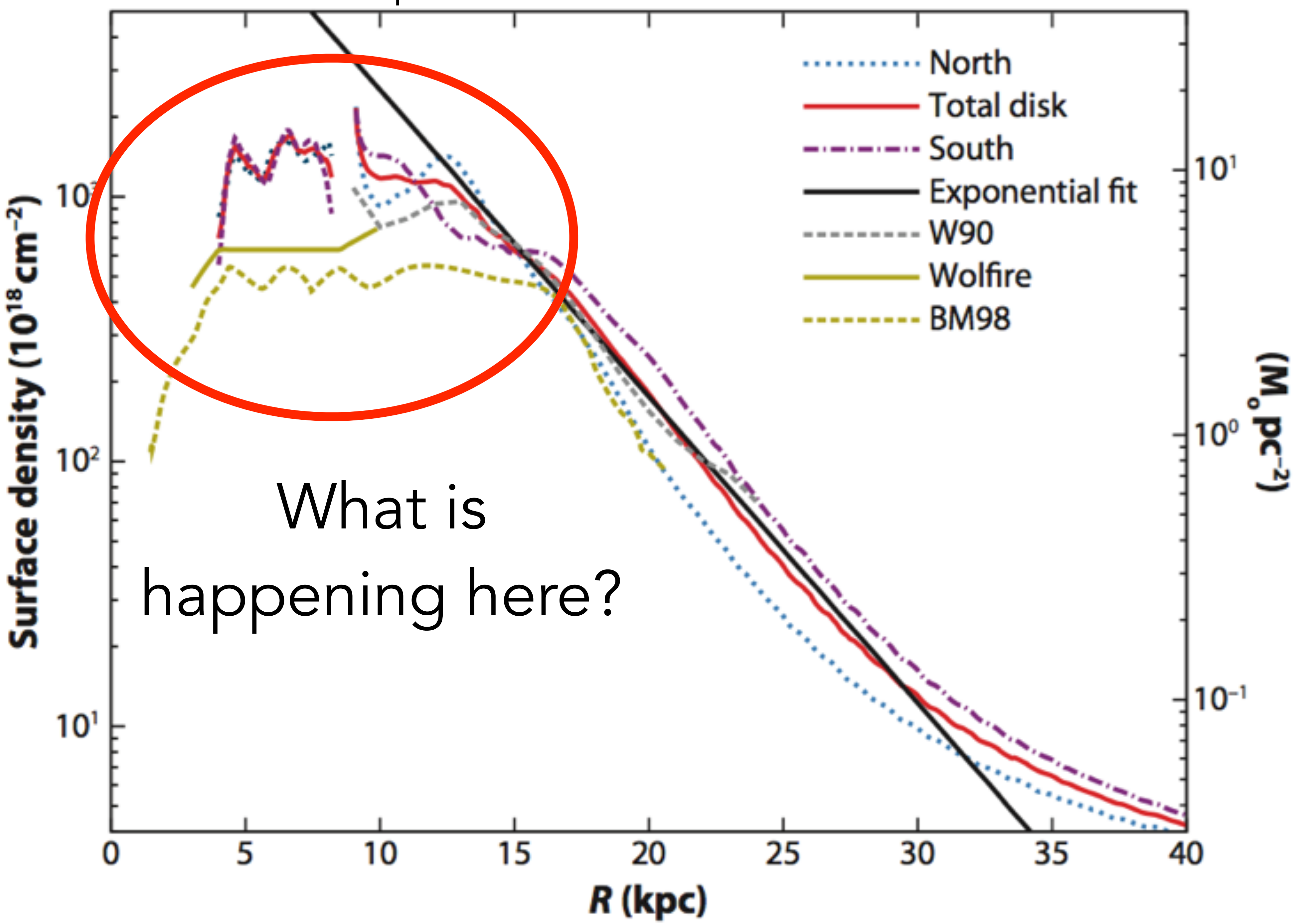

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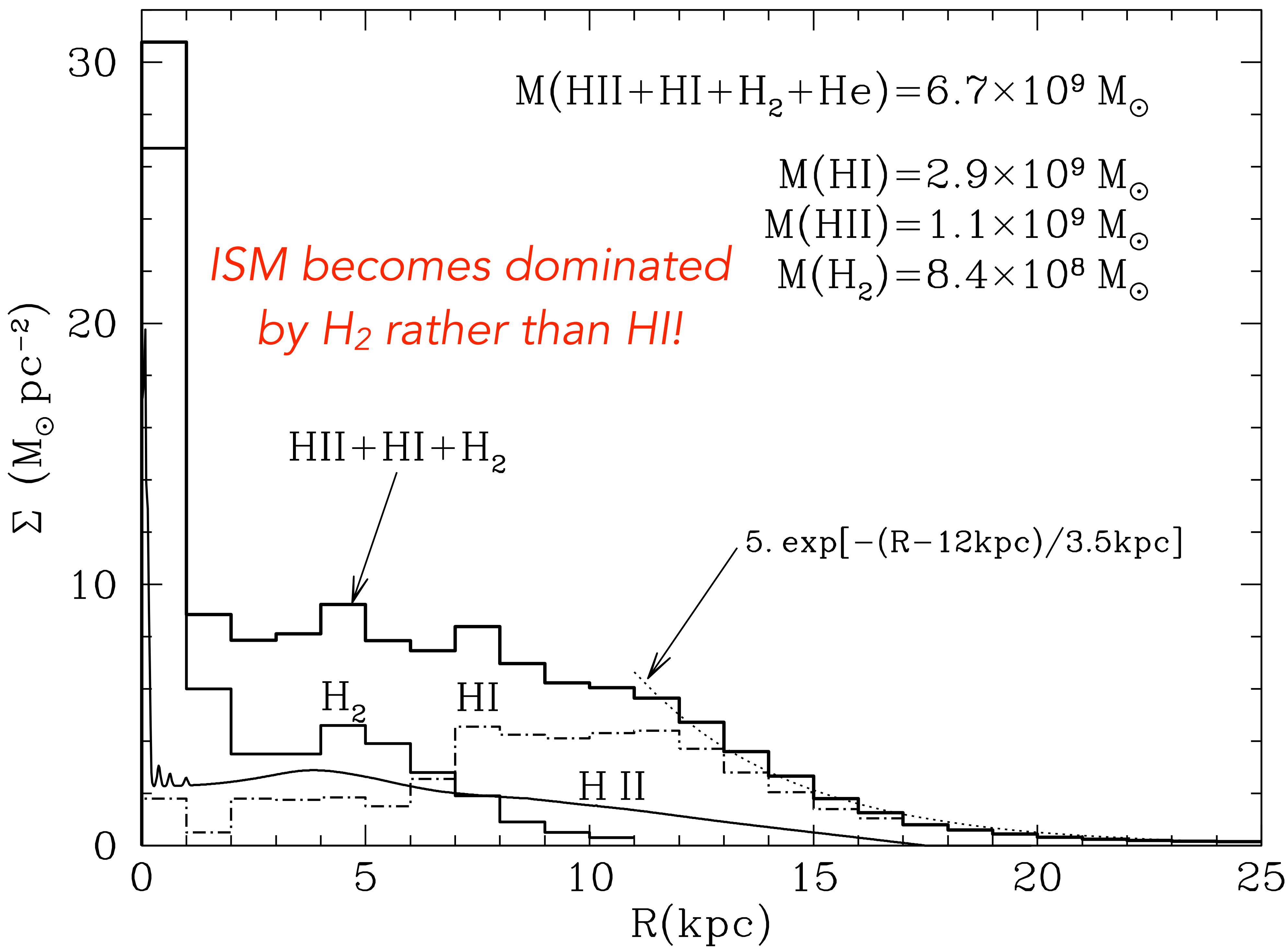
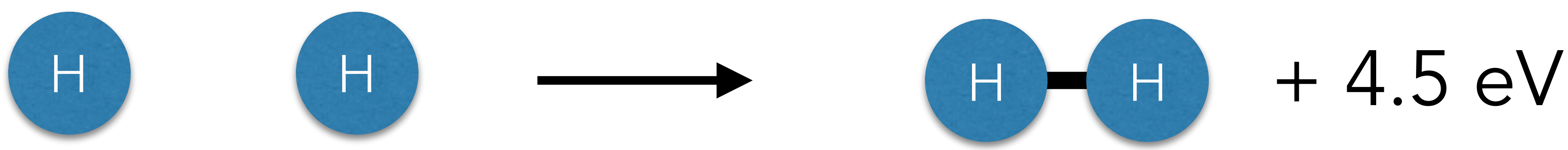
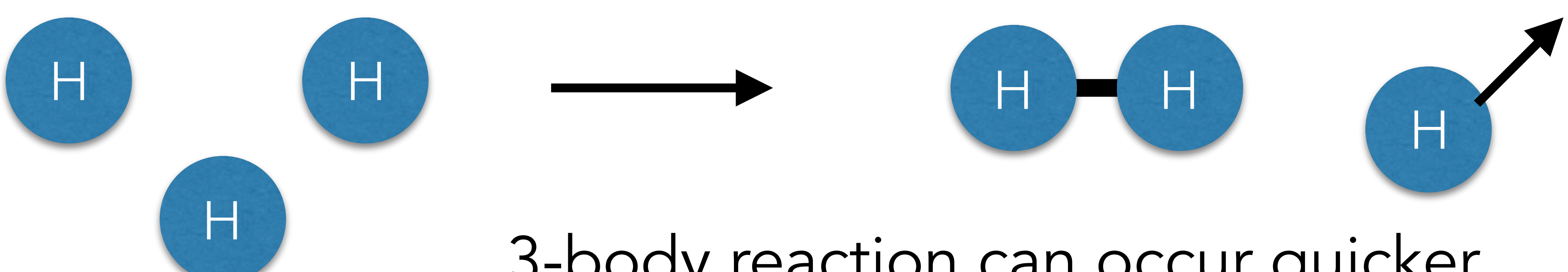


Forming H_2

Formation of H_2 by gas-phase reactions is slow



no effective way to carry away 4.5 eV worth of binding energy when two H bond, no dipole moment
negligible rate for this reaction

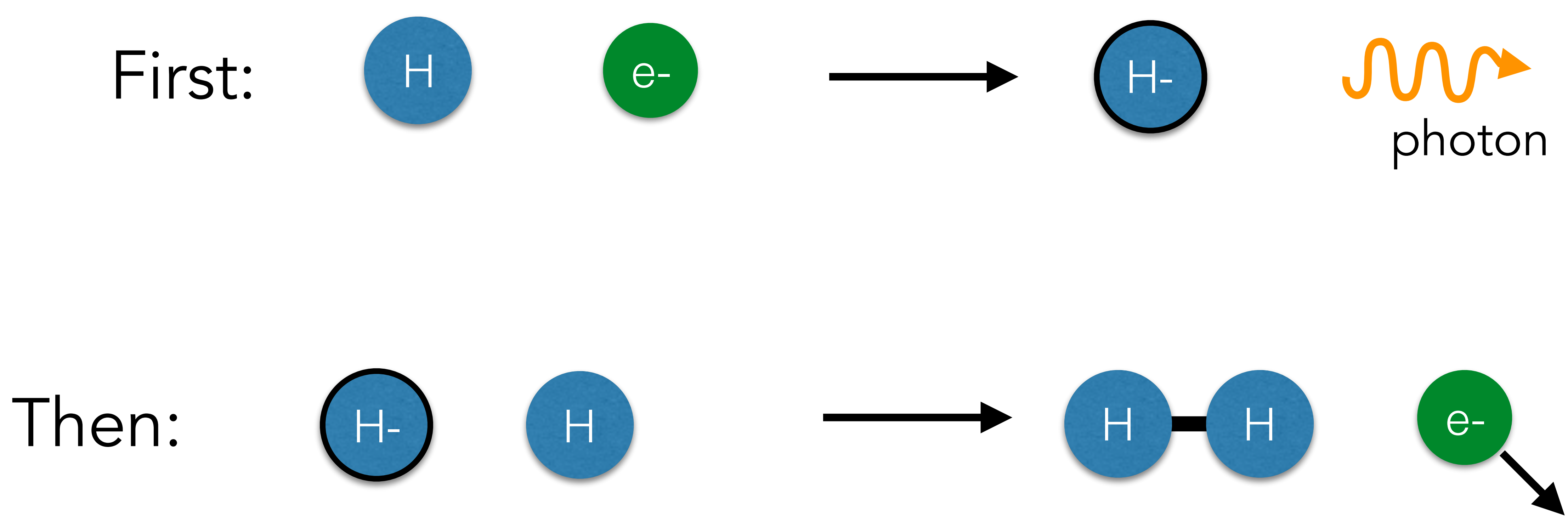


3-body reaction can occur quicker
but this is still very slow

Forming H_2

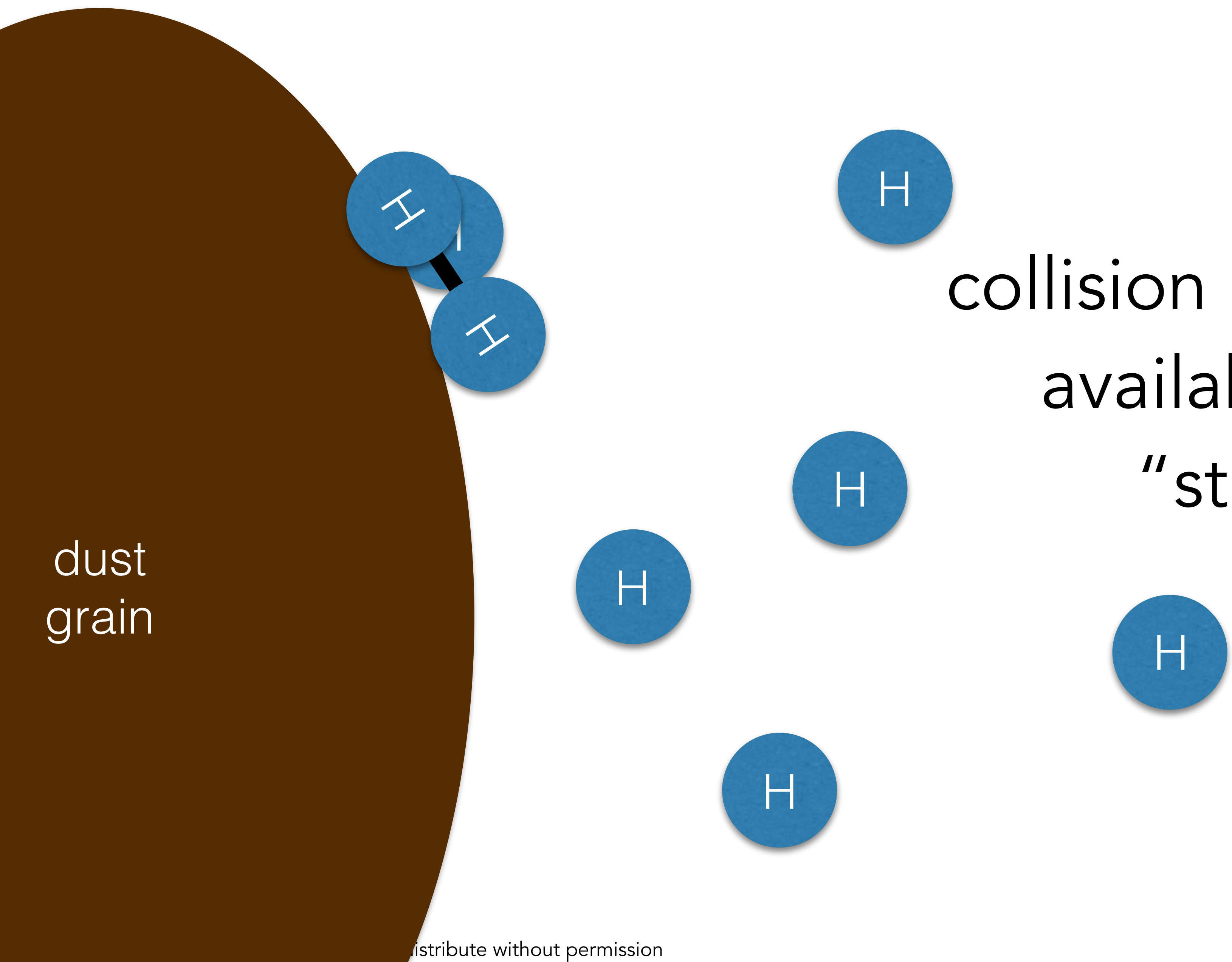
Formation of H_2 by gas-phase reactions is slow

Fastest gas-phase route is "associative attachment"



Forming H_2

Grain Surface H_2 formation is much faster if there is dust.



Depends on:
collision rate of H with grain (n,T)
available grain surface area
"sticking" probability

Forming H_2

Grain Surface H_2 formation is much faster if there is dust.

$$\left(\frac{dn(\mathrm{H_2})}{dt}\right)_{\mathrm{gr}} = R_{\mathrm{gr}}\ n_H\ n(H)$$

density of H atoms

total density of hydrogen in any form

$$R_{\mathrm{gr}} = \frac{1}{2}\left(\frac{8kT}{\pi m_H}\right)^{1/2} \langle\epsilon_{\mathrm{gr}}\rangle \Sigma_{\mathrm{gr}}$$

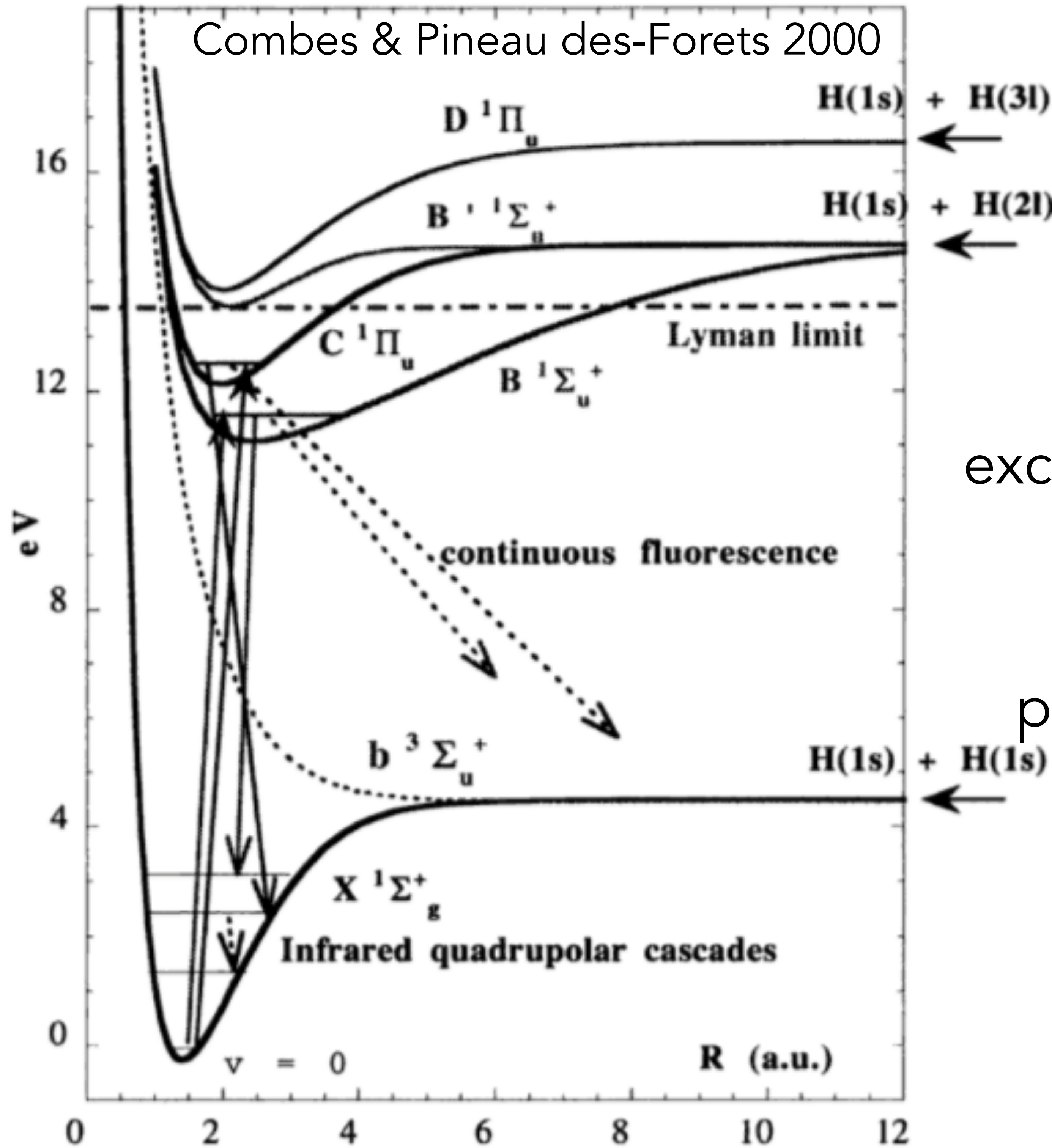
v_{thermal}

average "sticking" coeff for grain pop

Grain surface area

$$\Sigma_{\mathrm{gr}} \equiv \frac{1}{n_H}\int \pi a^2 \frac{dn_{\mathrm{gr}}}{da}da$$

Photodissociation of H_2



Combes & Pineau des-Forets 2000

After H_2 absorbs a UV photon from ground to one of the excited levels (Lyman-Werner bands) has ~85% probability of radiative decay, ~15% probability of photo-dissociating

Lyman band = ground -> B
Werner band = ground -> C

Photodissociation of H_2

$$\zeta_{\rm diss} = \sum_u \zeta_{l\rightarrow u}\, p_{{\rm diss},u}$$

dissociation rate

sum of rates for transitions from ground state to excited states

probability of photo-dissociation from upper state

depends on quantum mechanics and radiation field intensity at relevant wavelengths

H_2 Abundance

In steady state:

photo-dissociation

$$\varsigma_{\rm diss} n(H_2) = R_{\rm gr} n_H n(H)$$

formation on dust grains

For CNM conditions this is pretty small:

$$\frac{n(H_2)}{n_H} \approx 1.8 \times 10^{-5} \left(\frac{n(H)}{30{\rm cm}^{-3}}\right) \left(\frac{R_{\rm gr}}{3 \times 10^{-17}{\rm cm}^3{\rm s}^{-1}}\right) \left(\frac{\varsigma_{\rm diss}}{5 \times 10^{-11}{\rm s}^{-1}}\right)^{-1}$$

But we have left out an important component: shielding

H_2 Abundance

Sternberg et al. 2014

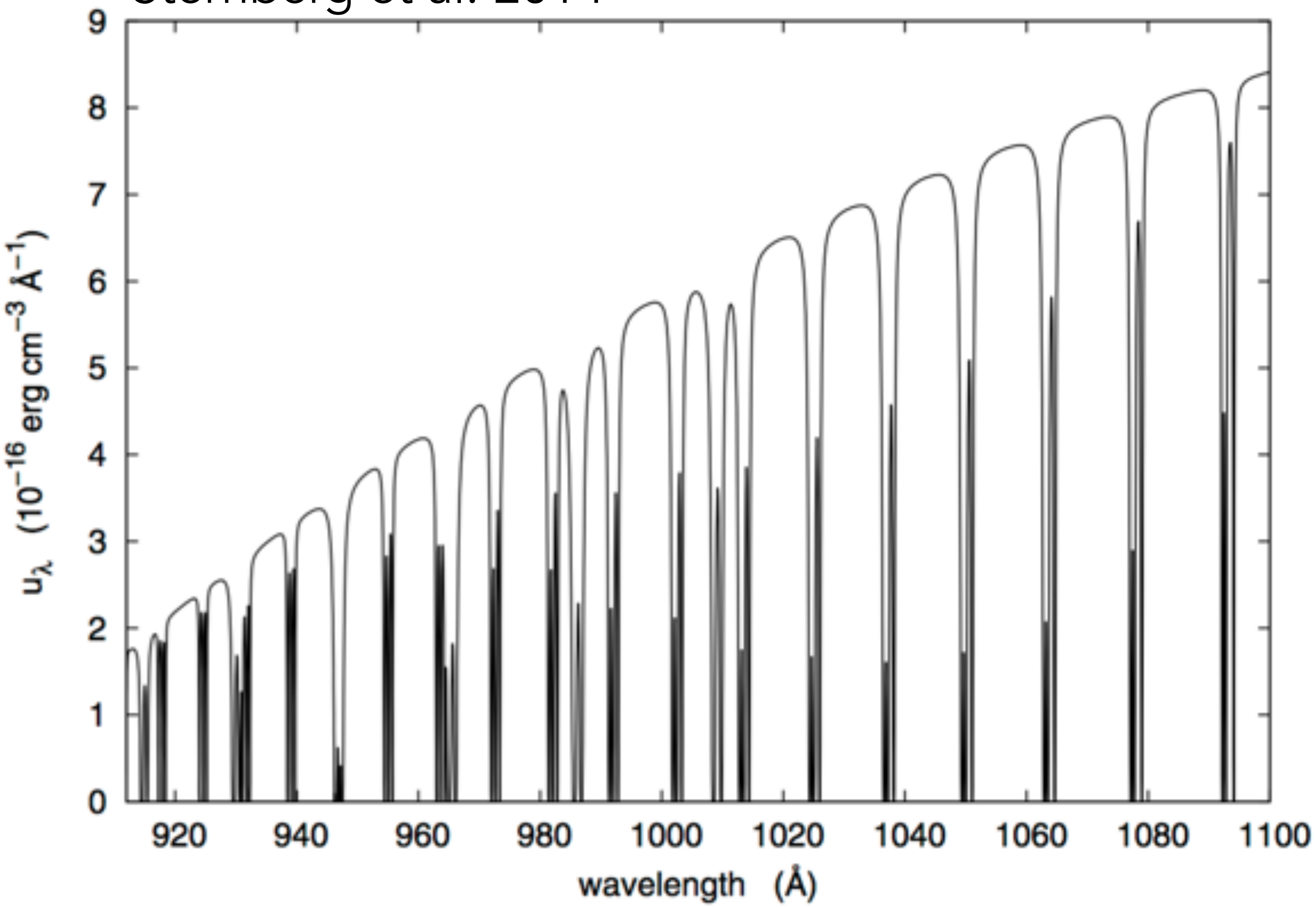
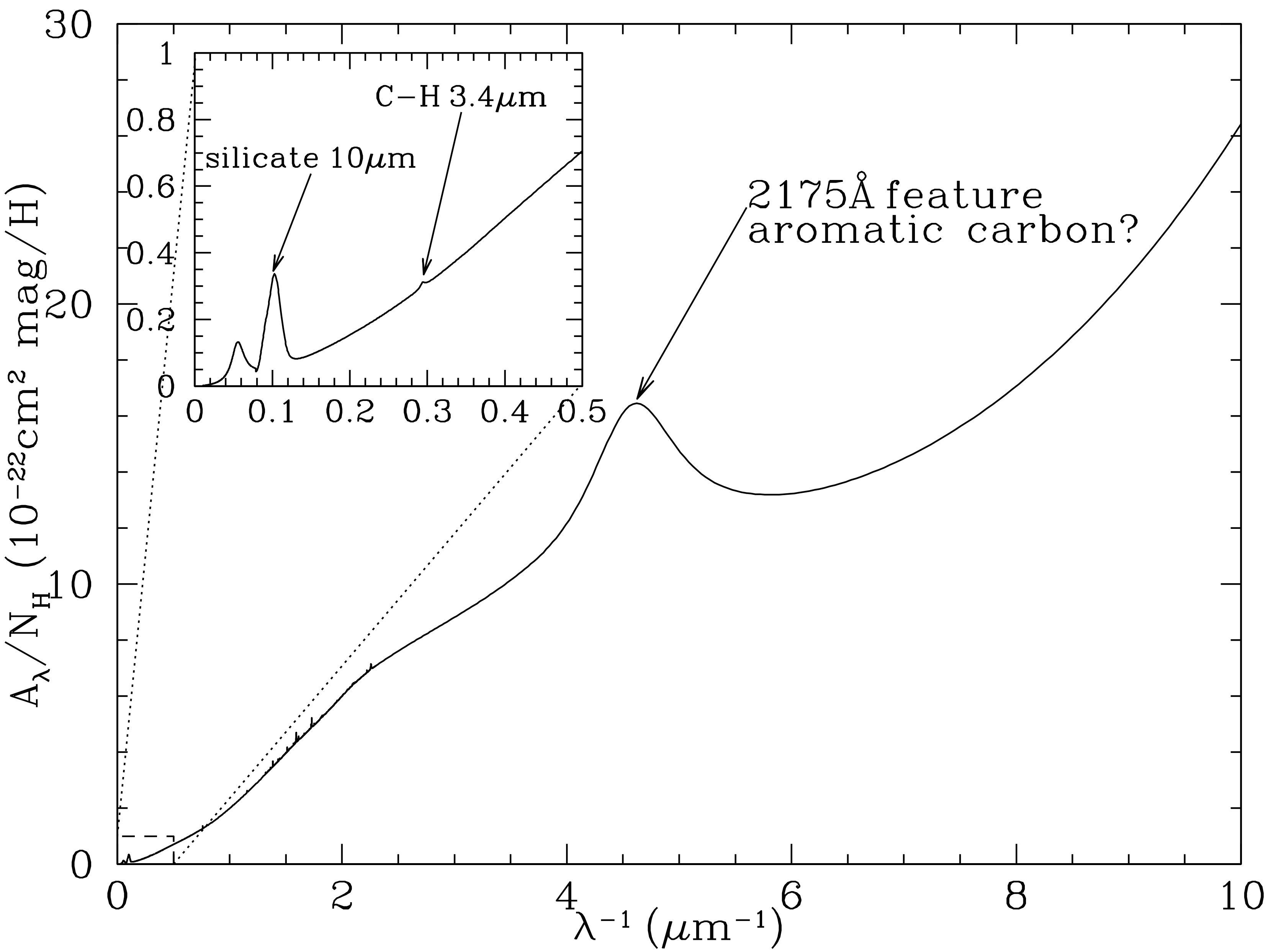


Figure 2. Absorbed far-UV spectrum showing partially overlapping Lyman–Werner band absorption lines, for beamed radiation into a cloud, at a total hydrogen gas column density of 3.74×10^{20} cm^{-2}, for a free-space radiation intensity $I_{UV} = 35.5$, gas density $n = 10^3$ cm^{-3}, and metallicity $Z' = 1$ $(\alpha G/2 = 1)$.

H_2 Lyman-Werner bands can become optically thick and shield interior H_2 from being dissociated.

H_2 Abundance



At UV wavelengths even small A_V corresponds to substantial amounts of UV extinction.

H_2 Abundance

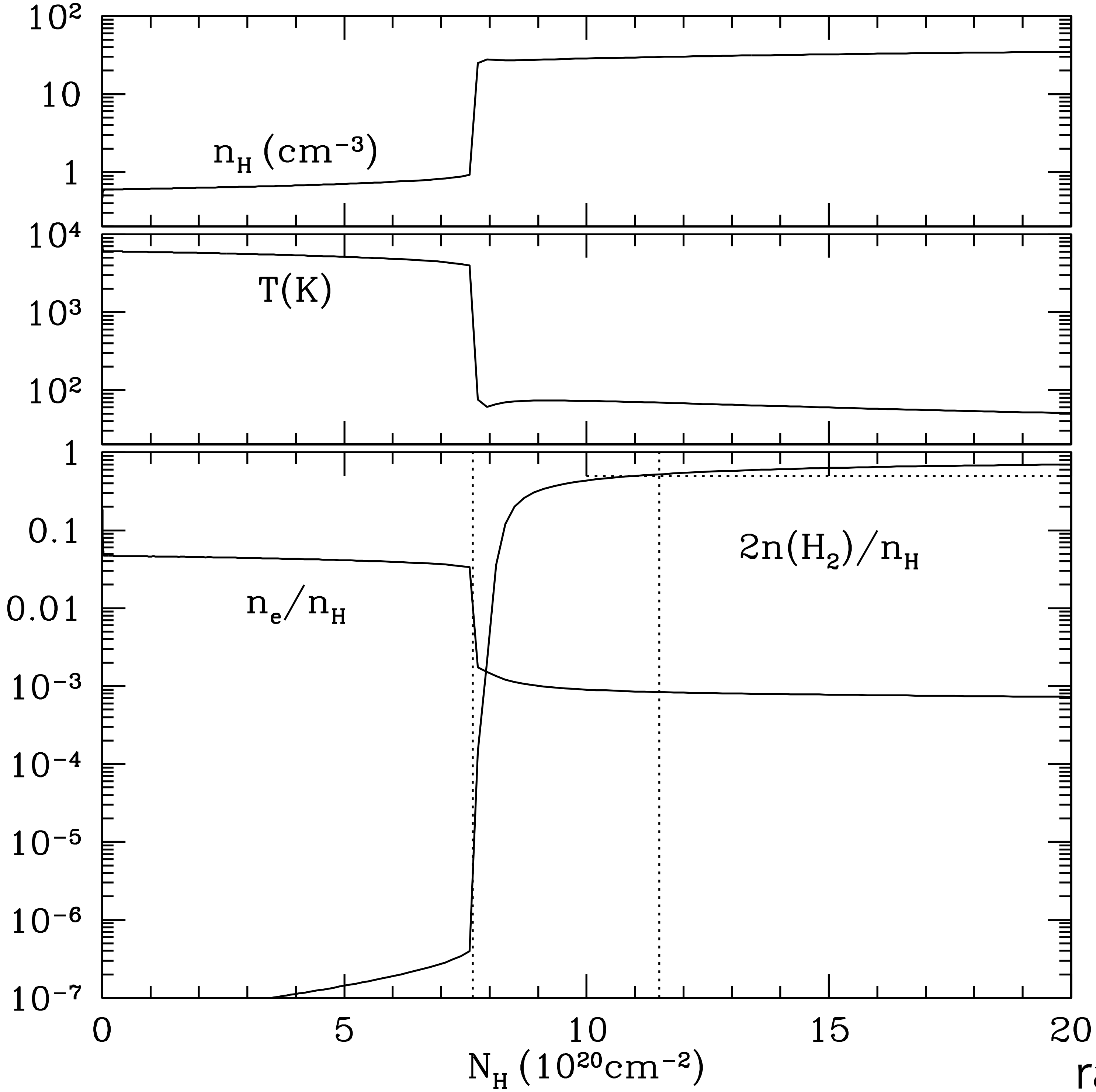
$$\varsigma_{\rm diss} \approx \varsigma_{\rm diss,0} \; f_{\rm shield} \; e^{-\tau_{d,1000}}$$

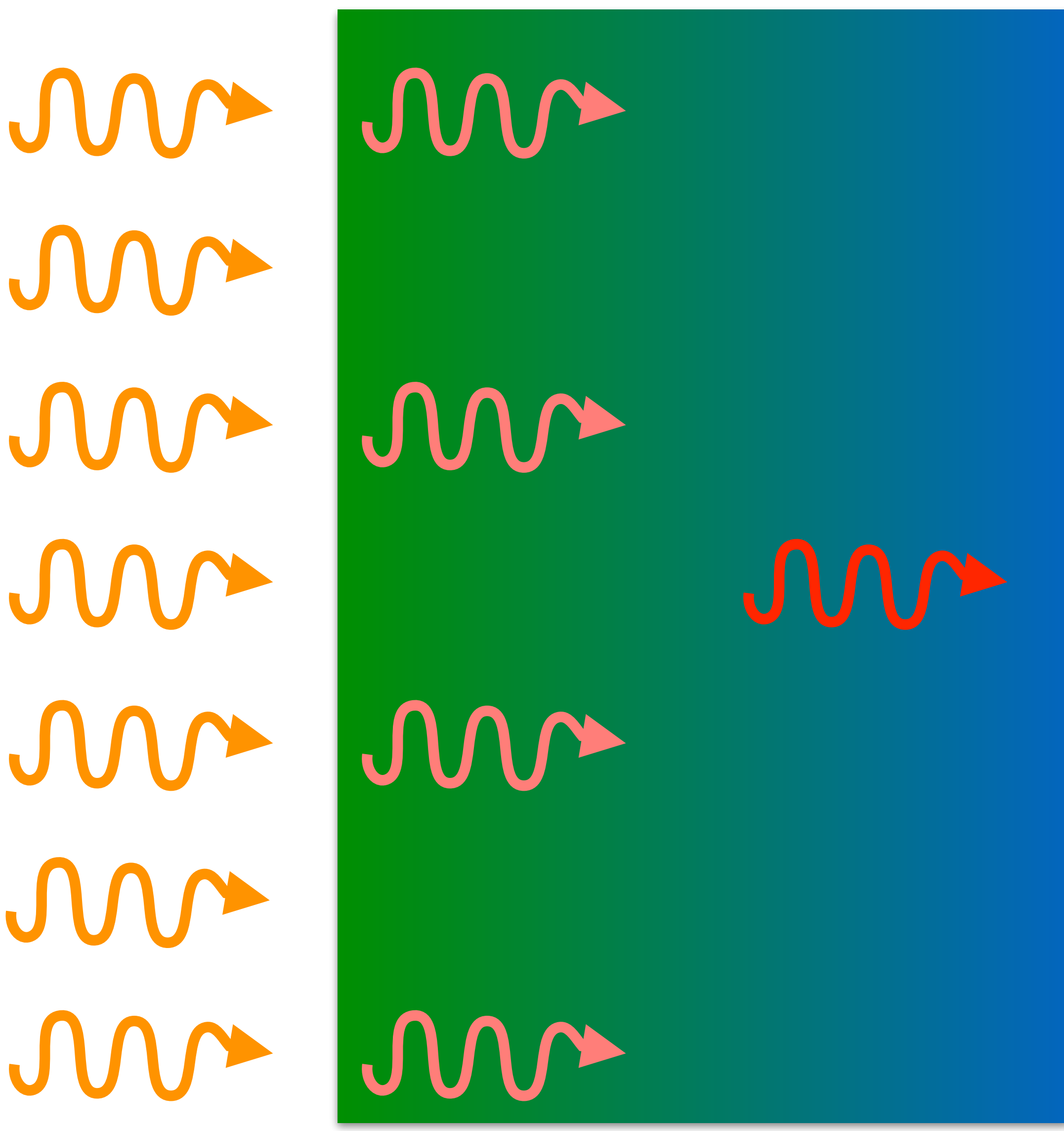
dissociation rate with no shielding

self-shielding factor

dust extinction at 1000 Å

H_2 Abundance





external radiation field

plane-parallel distribution of gas

Photodissociation Regions

Very general term, can refer to anywhere that far-UV (<13.6 eV) photons play key role in chemistry, ionization, etc.