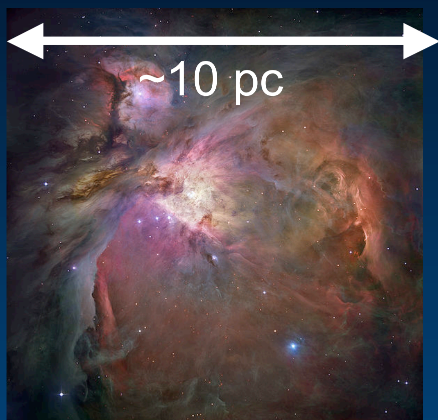
Galaxy Simulators
Star Formation for ~~Dummies~~
^

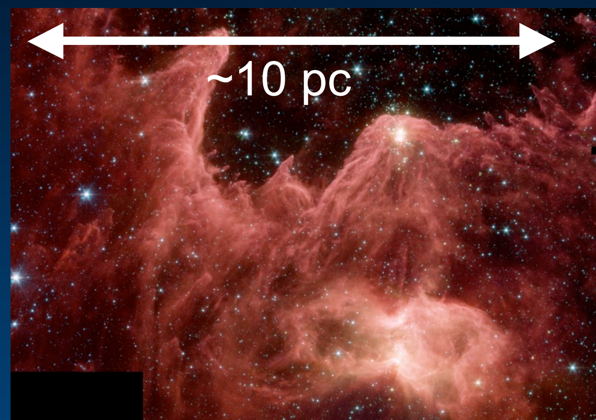
Mark Krumholz
UC Santa Cruz

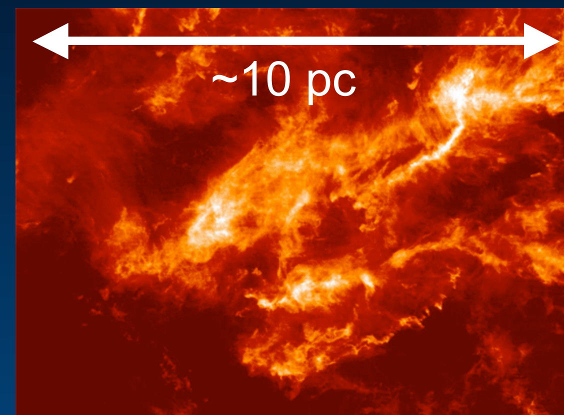
HIPACC Summer School

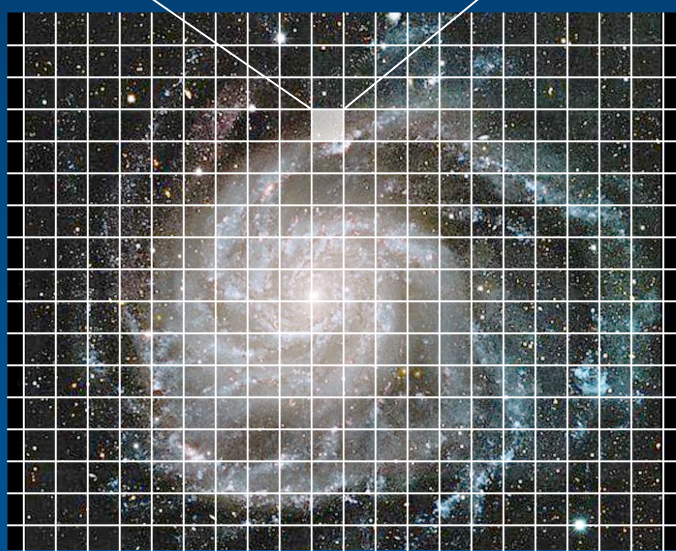
August 6, 2010

The Challenge of Star Formation









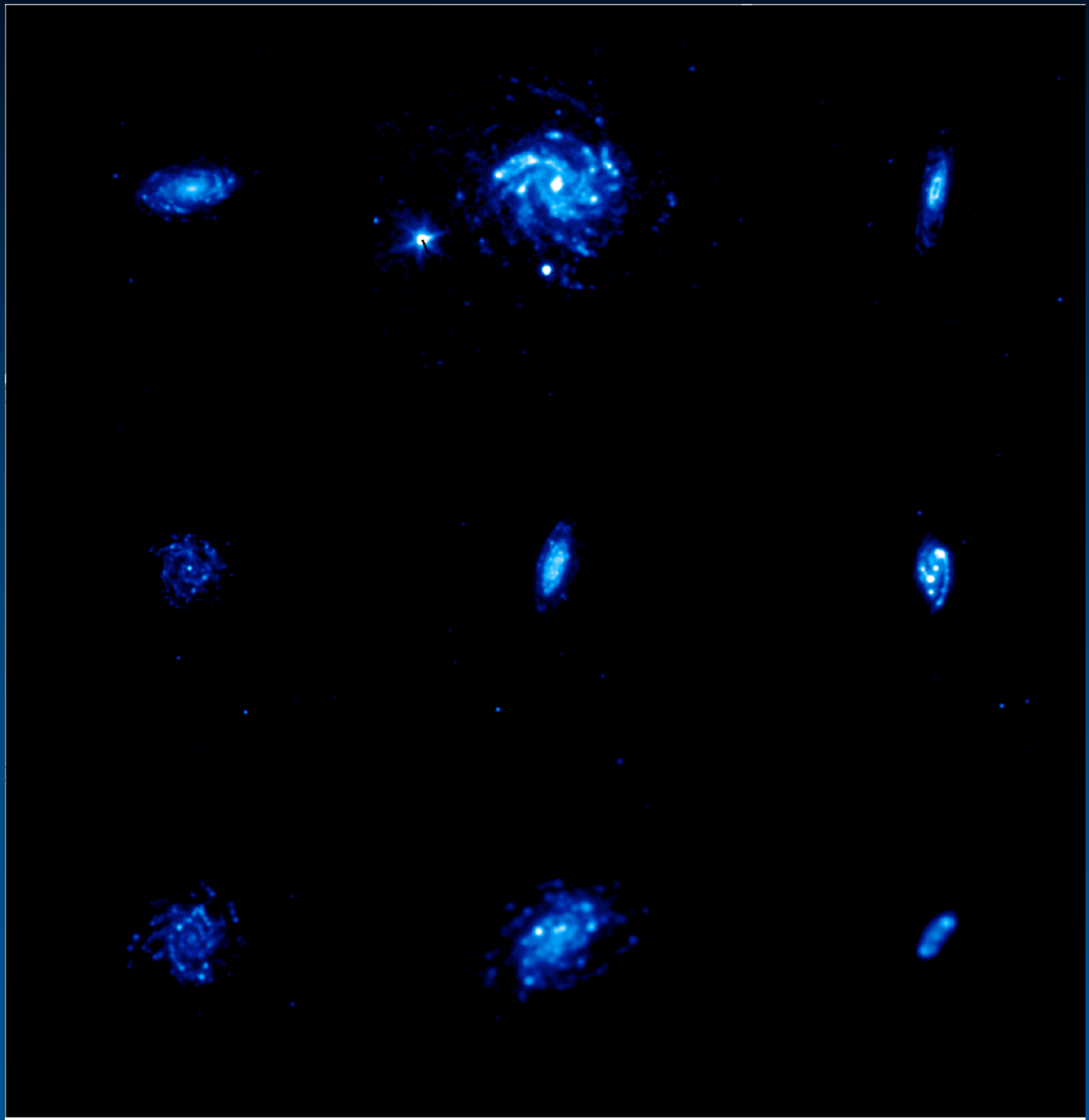
Like stars, star formation involves impossibly small length / time scales
⇒ must rely on subgrid models

Stellar models are very reliable; star formation models are not…

Subgrid Models that are Less Bad

- Stars form only in molecular gas; where the gas becomes molecular depends on metallicity
- Even in molecular gas, star formation is very slow; gas depletion times ~ t_H / 10
- Star formation feedback means more than just supernovae

A Story of HI and H_2

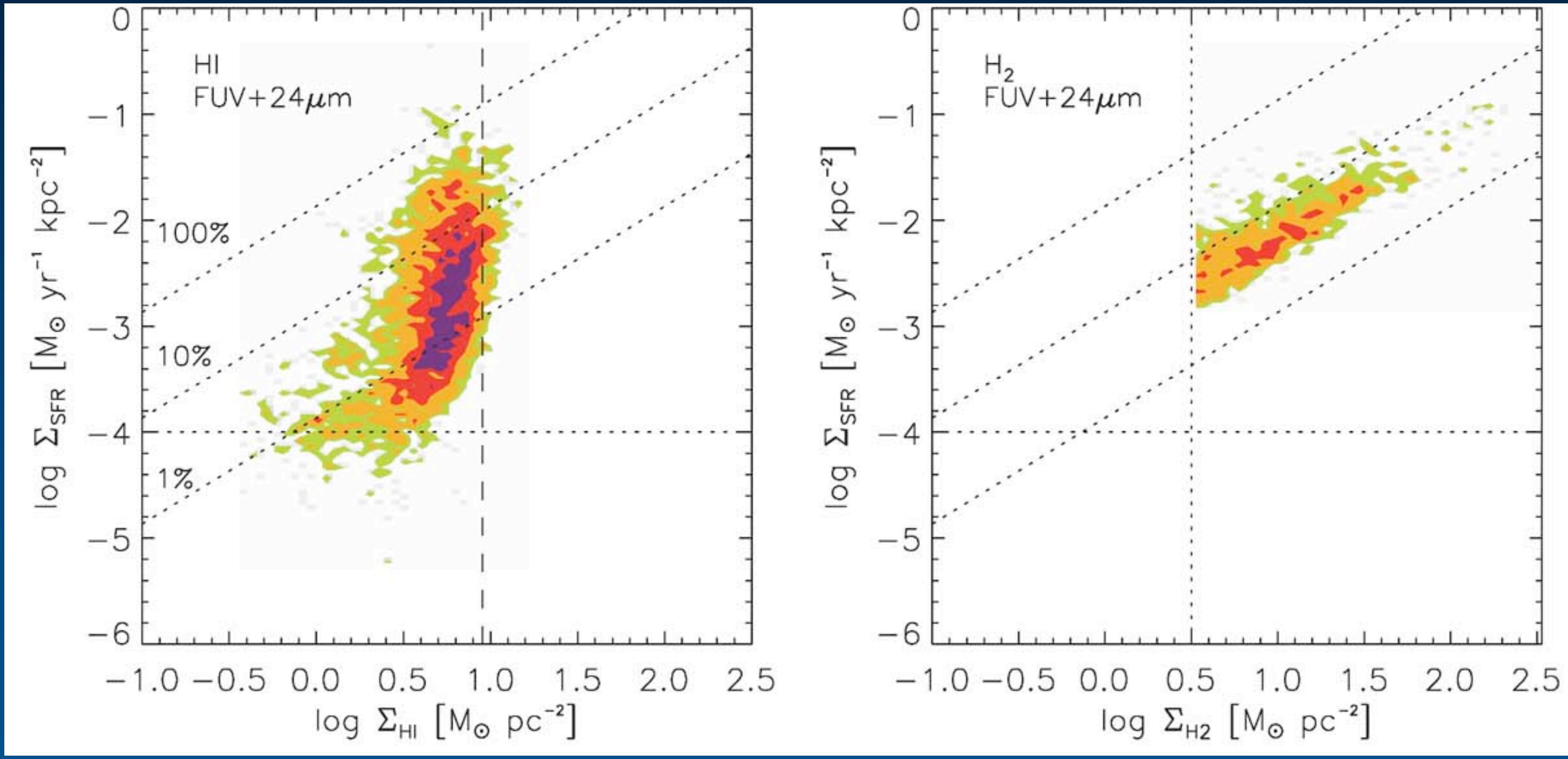


SFR distributions from 24 μm SINGS + GALEX

SINGS + GALEX + THINGS + SONG (animation borrowed from N. Gnedin)

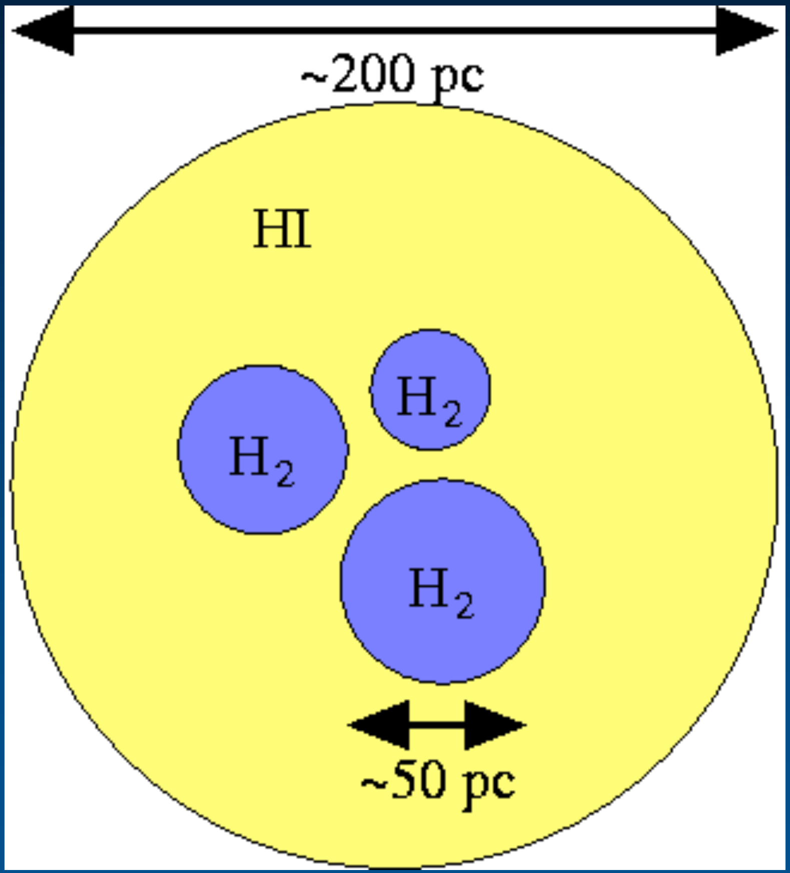
The SFR in HI and H_2

(Bigiel et al. 2008)



What Sets the HI / H_2 Transition?

- Molecules reside in giant molecular clouds (GMCs) that are the inner parts of atomic-molecular complexes
- The outer parts are dissociated by interstellar Lyman-Werner photons
- Goal: compute HI and H_2 mass fractions



How to Do this Numerically

(Pelupessy+ 2006, 2009, 2010; Robertson & Kravtsov 2008; Gnedin+ 2009, 2010)

- Update chemical state following

$$\frac{dn_{\mathrm{H_2}}}{dt} = -\frac{dn_{\mathrm{HI}}}{dt} = n_{\mathrm{HI}} n \mathcal{R} - \int d\Omega \int d\nu \, \sigma_{\mathrm{H_2}} f_{\mathrm{diss}} I_\nu / h\nu$$

- Rate coefficient $\mathcal{R}$ ∝ dust / gas ratio ∝ Z
- Approximate radiation field I_ν ∝ local SFR, or using an approximate radiative transfer method
- Star formation rate proportional to molecular mass / density only

An Analytic Model for HI / H_2 Balance

(Krumholz, McKee, & Tumlinson 2008, 2009; McKee & Krumholz 2010)

Approximate the system as being in approximate chemical equilibrium.

Formation on grains = Photodissociation

$$
\begin{aligned}
n_{\mathrm{HI}} n \mathcal{R} &= n_{\mathrm{H_2}} \int d\Omega \int d\nu \, \sigma_{\mathrm{H_2}} f_{\mathrm{diss}} I_\nu / (h\nu) \\
\hat{e} \cdot \nabla I_\nu &= -(n_{\mathrm{H_2}} \sigma_{\mathrm{H_2}} + n \sigma_{\mathrm{d}}) I_\nu
\end{aligned}
$$

Decrease in radiation intensity = Absorptions by H_2 molecules + dust grains

Idealized problem: spherical cloud of radius R, density n, dust opacity σ_d, H_2 formation rate coefficient $\mathcal{R}$, unshielded radiation field with LW photon number density E_0^*, find fraction of mass in HI and H_2.

Calculating Molecular Fractions

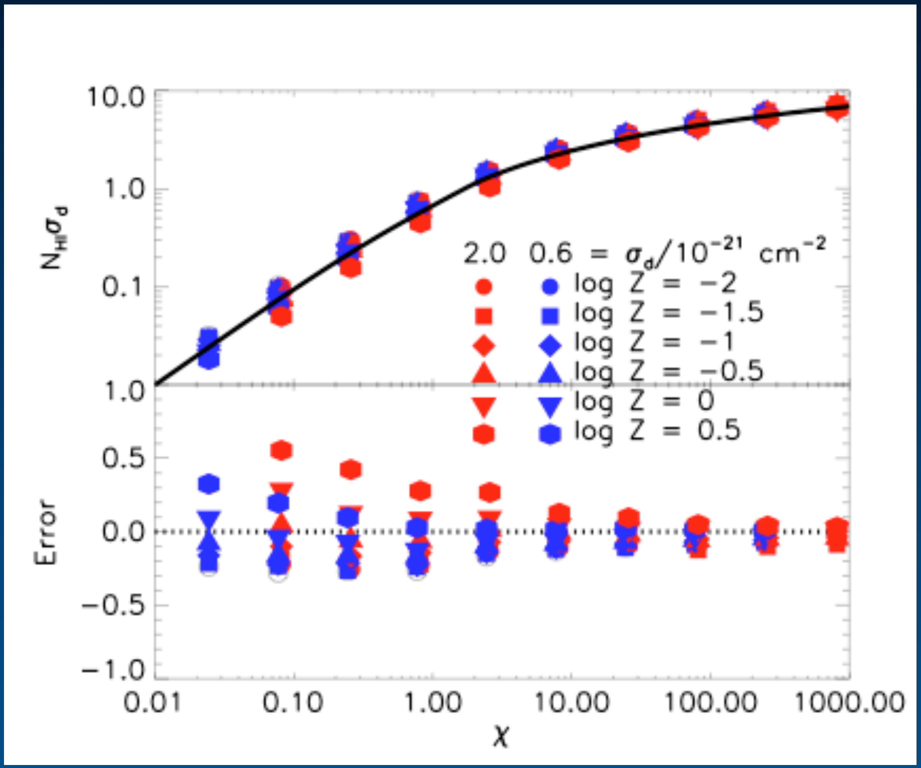
To good approximation, solution only depends on two numbers:

$$\tau_R = n\sigma_d R$$

$$\chi = \frac{f_{diss}\sigma_d E_0^*}{n\mathcal{R}}$$

A semi-analytic solution can be given from these parameters.

τ_R depends only on galaxy Σ, $Z \Rightarrow$ can be measured directly



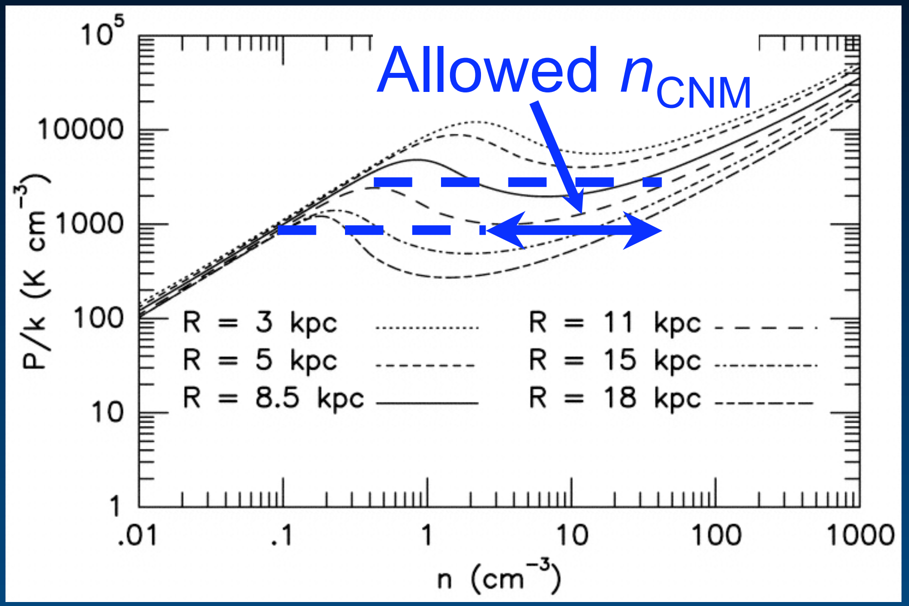
Analytic solution for location of HI / H_2 transition vs. exact numerical result

Shielding Layers in Galaxies

(Krumholz, McKee, & Tumlinson 2009)

What is $\chi \propto (\sigma_d / \mathcal{R})\ (E_0^* / n)$?

- Dust opacity σ_d and H_2 formation rate $\mathcal{R}$ both $\propto$ Z, so $\sigma_d / \mathcal{R}$ ~ const
- CNM dominates shielding, so n is the CNM density



FGH curves for MW (Wolfire et al. 2003)

- CNM density set by pressure balance with WNM, and $n_{CNM} \propto E_0^*$, with weak Z dependence.

$\Rightarrow$ $\chi \propto (\sigma_d / \mathcal{R})\ (E_0^* / n)$ ~ 1 in all galaxies!

Predictions for H_2 Content

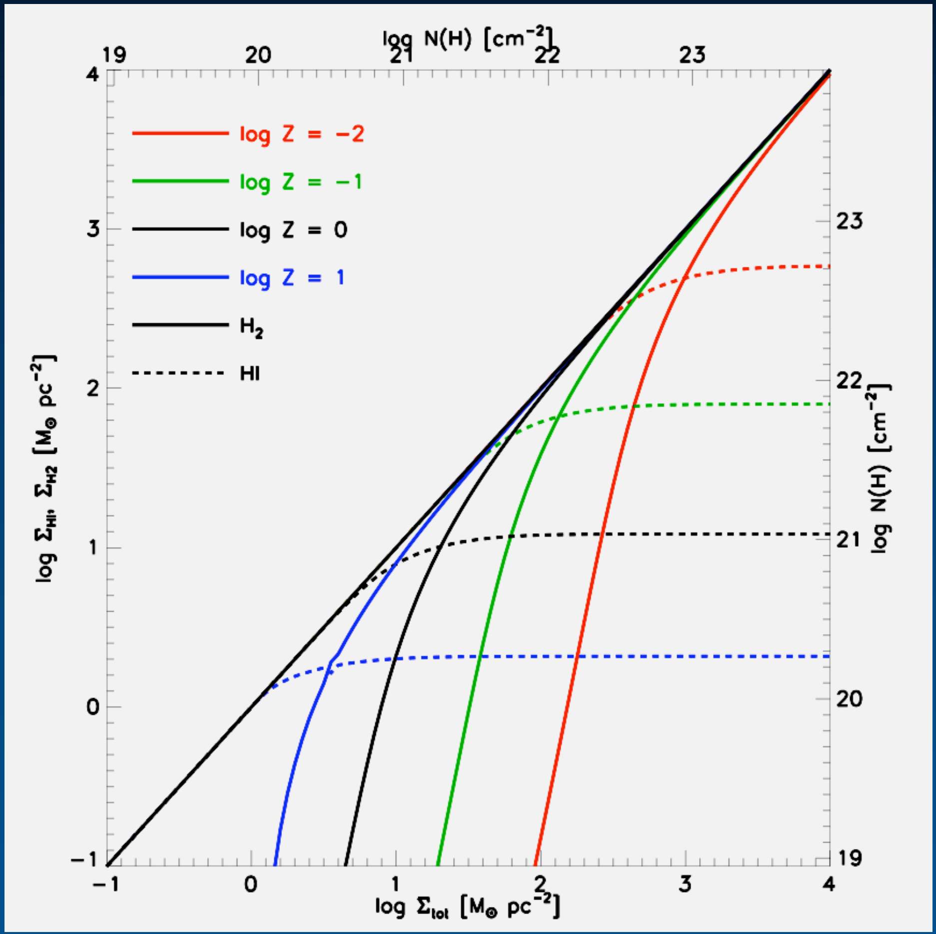
- Bottom line

$$f_{H_2} \approx 1 - \frac{3}{4}\left(\frac{s}{1+0.25s}\right)$$

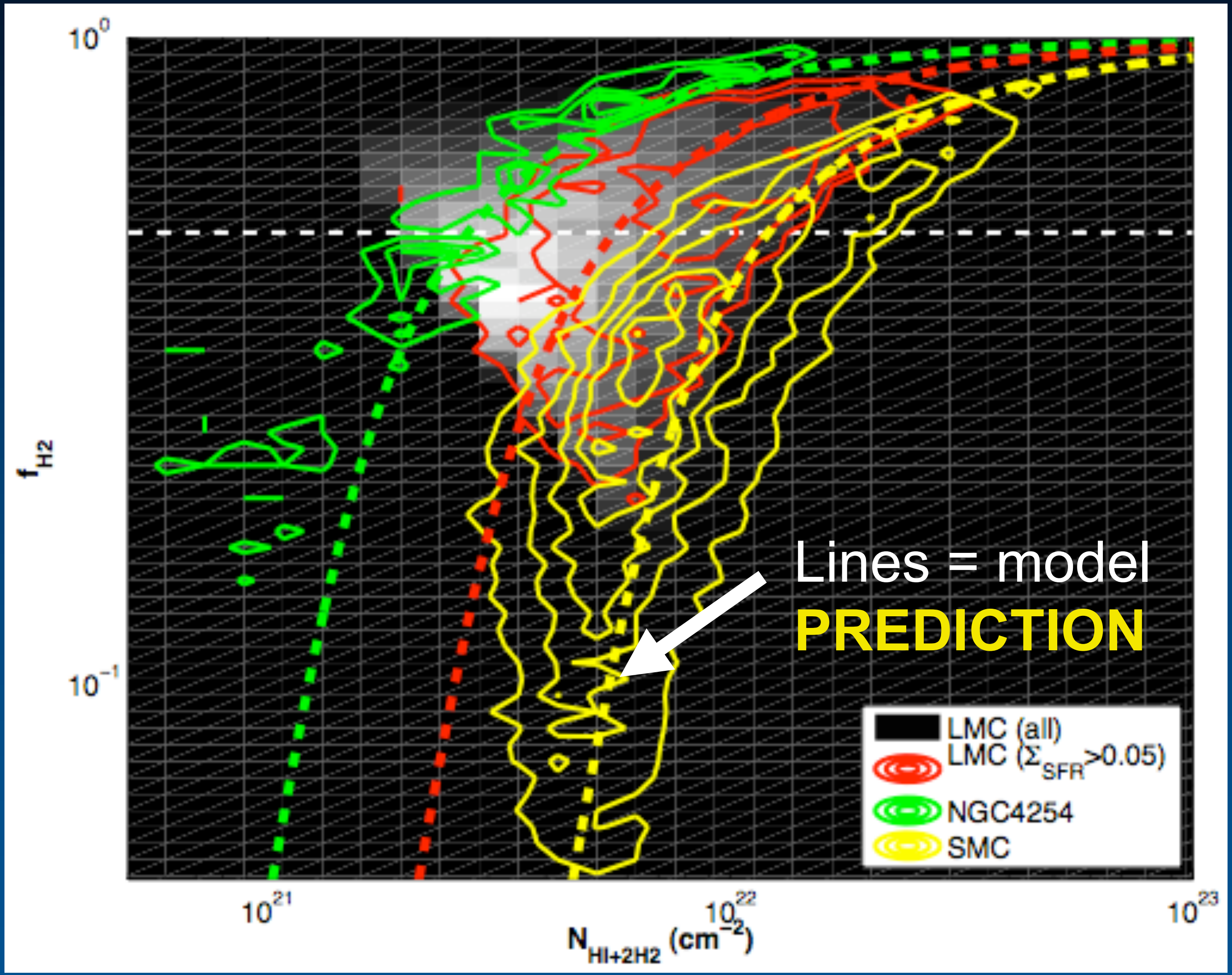
$$s \approx \frac{\ln(1+0.6\chi+0.01\chi^2)}{0.04\left(\frac{Z}{Z_\odot}\right)\left(\frac{\Sigma}{M_\odot\ \mathrm{pc}^{-2}}\right)}$$

$$\chi \approx 3.1\frac{1+3.1\left(\frac{Z}{Z_\odot}\right)^{0.365}}{4.1}$$

- Qualitative effect: f_{H2} goes from ~0 to ~1 when $\Sigma Z \sim 10\ M_\odot\ \mathrm{pc}^{-2}$



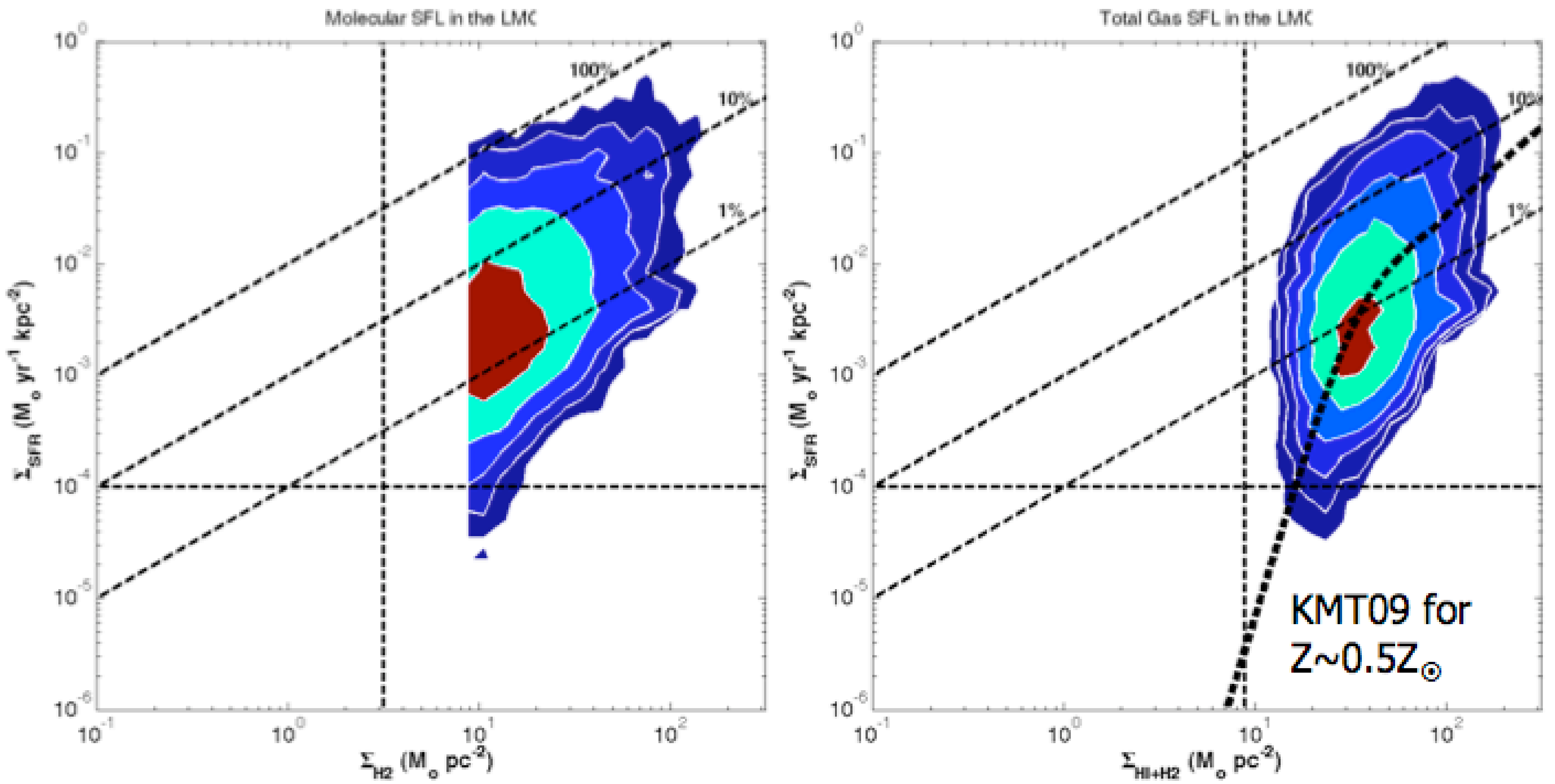
Checking the Model



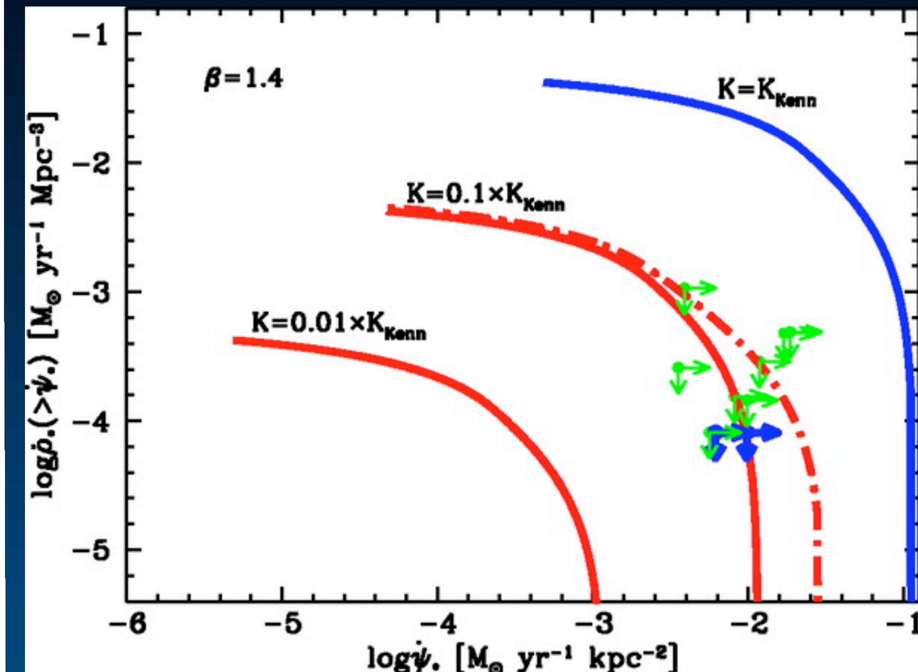
Molecular fractions in nearby galaxies, with H_2 inferred from dust (Bolatto+ 2010, in preparation)

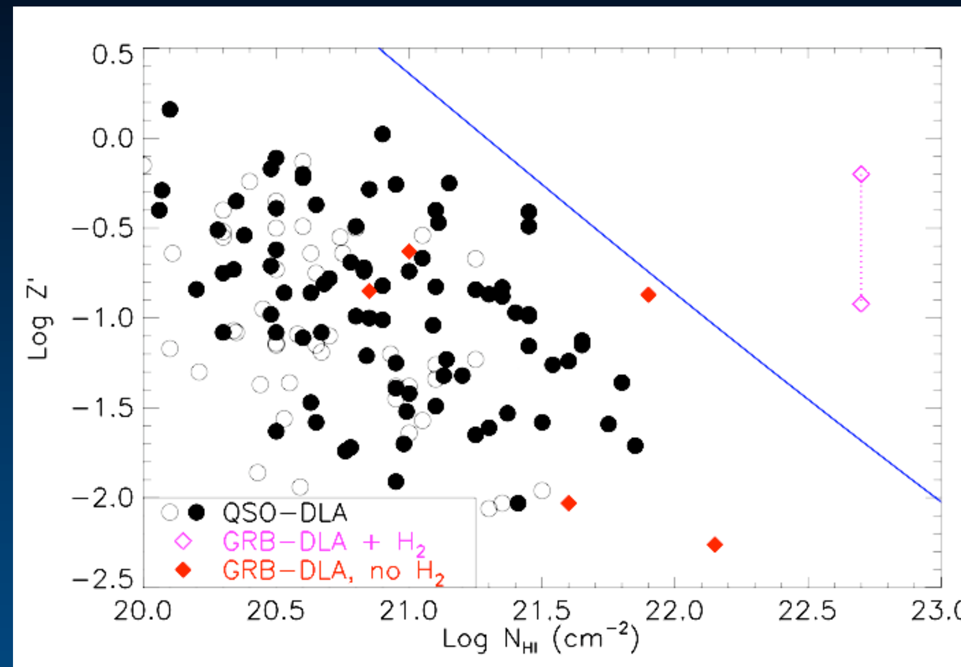
Why This Matters: SMC and LMC

(data from Bolatto+ 2010, in preparation)



Why This Matters: DLAs





Left: z ~ 3 galaxy density if the DLAs follow the Kennicutt (1998) SF law; data plus expectations (Wolfe & Chen 2006)

Right: DLA column density and metallicity distribution, plus line showing HI - H_2 transition (Krumholz+ 2009)

How to Use This: A Suggestion

- Must track metallicity
- Follow H_2 in non-equilibrium using chemical evolution (more accurate)
- Take $\Sigma \sim \rho_{HI}$ h, with scale height given by $h = \rho_{HI} / |\nabla\rho_{HI}|$, compute H_2 fraction from KMT equilibrium approximation (less accurate, but faster)

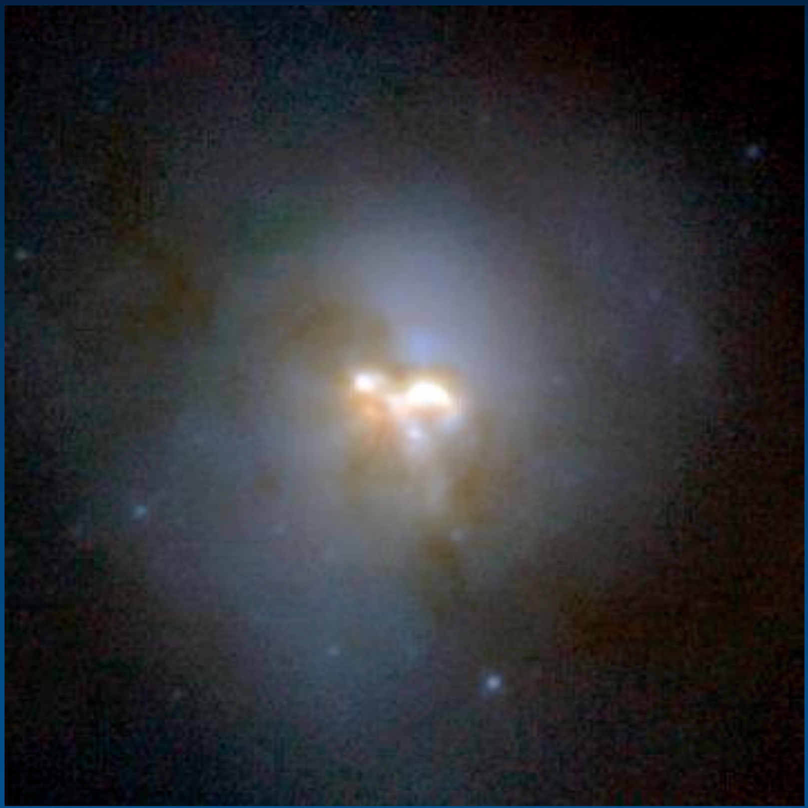
Star Formation is Slow…

(Zuckerman & Evans 1974; Rownd & Young 1999; Wong & Blitz 2002)

- The MW disk contains ~10^9 $M_\odot$ of gas in giant molecular clouds
- GMCs have n_H ~ 100 cm^{-3}, t_{ff} ~ 4 Myr
- If GMCs were collapsing, the SFR would be ~10^9 $M_\odot$ / 4 Myr = 250 $M_\odot$ / yr
- Observed SFR in MW is ~ 1 $M_\odot$ / yr, lower by a factor of ~100
- Numbers similar in nearby galaxies

…even in starbursts…

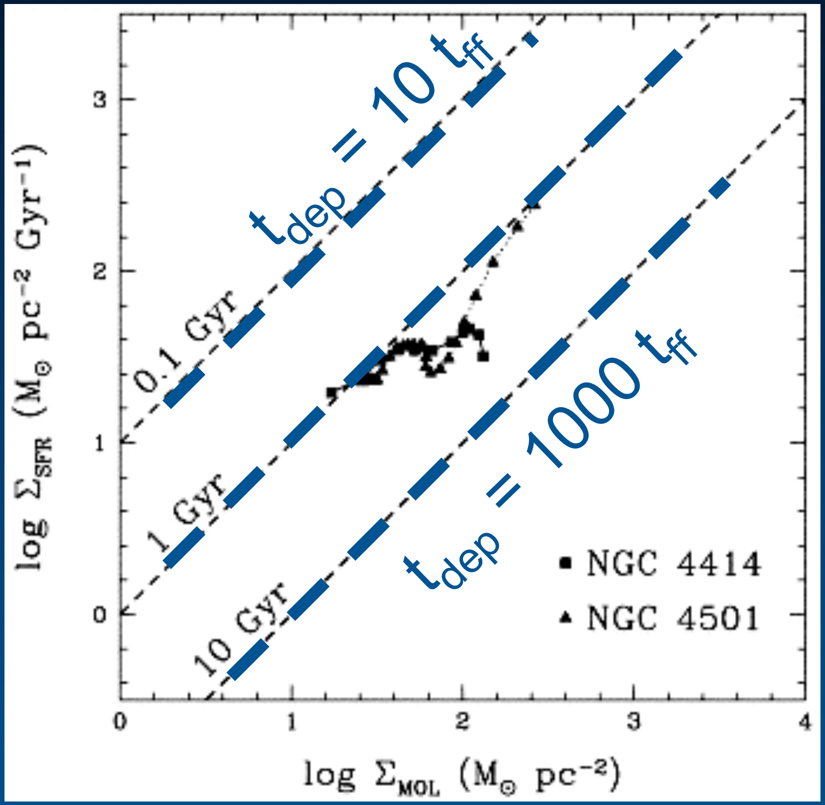
(Downes & Solomon 1998)

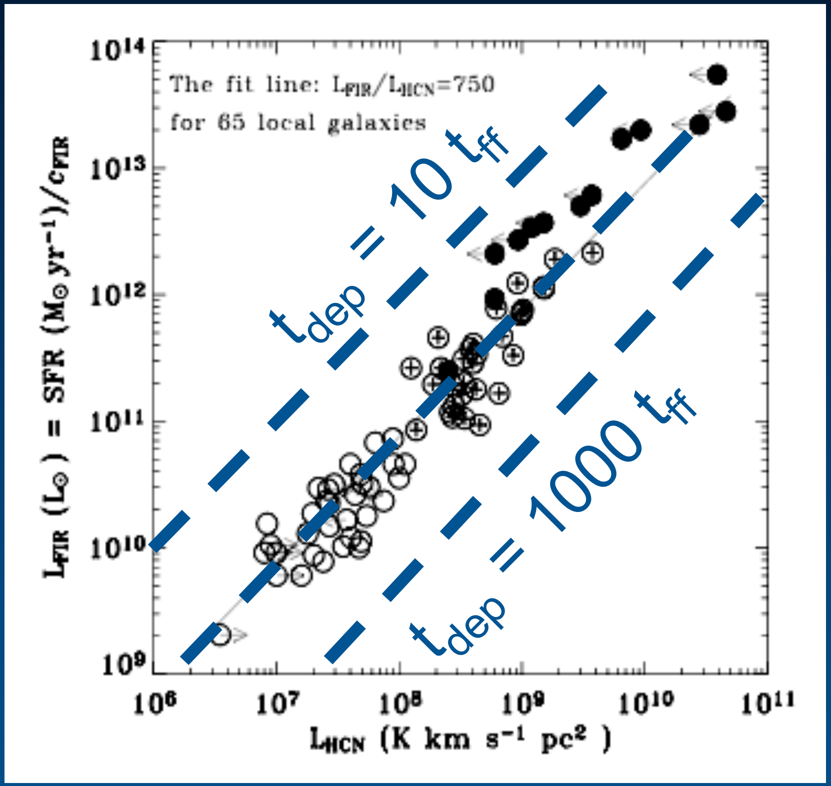


Arp 220 imaged by HST/NICMOS, Thompson et al. 1997

- Example: Arp 220
- ISM mass 2 x 10^9 $M_\odot$ in molecular gas
- ISM density 10^4 cm^{-3}, t_{ff} ~ 0.4 Myr
- Suggested SFR ~ 5000 $M_\odot$ / yr
- Actual SFR ~ 50 $M_\odot$ / yr : too small by factor of 100

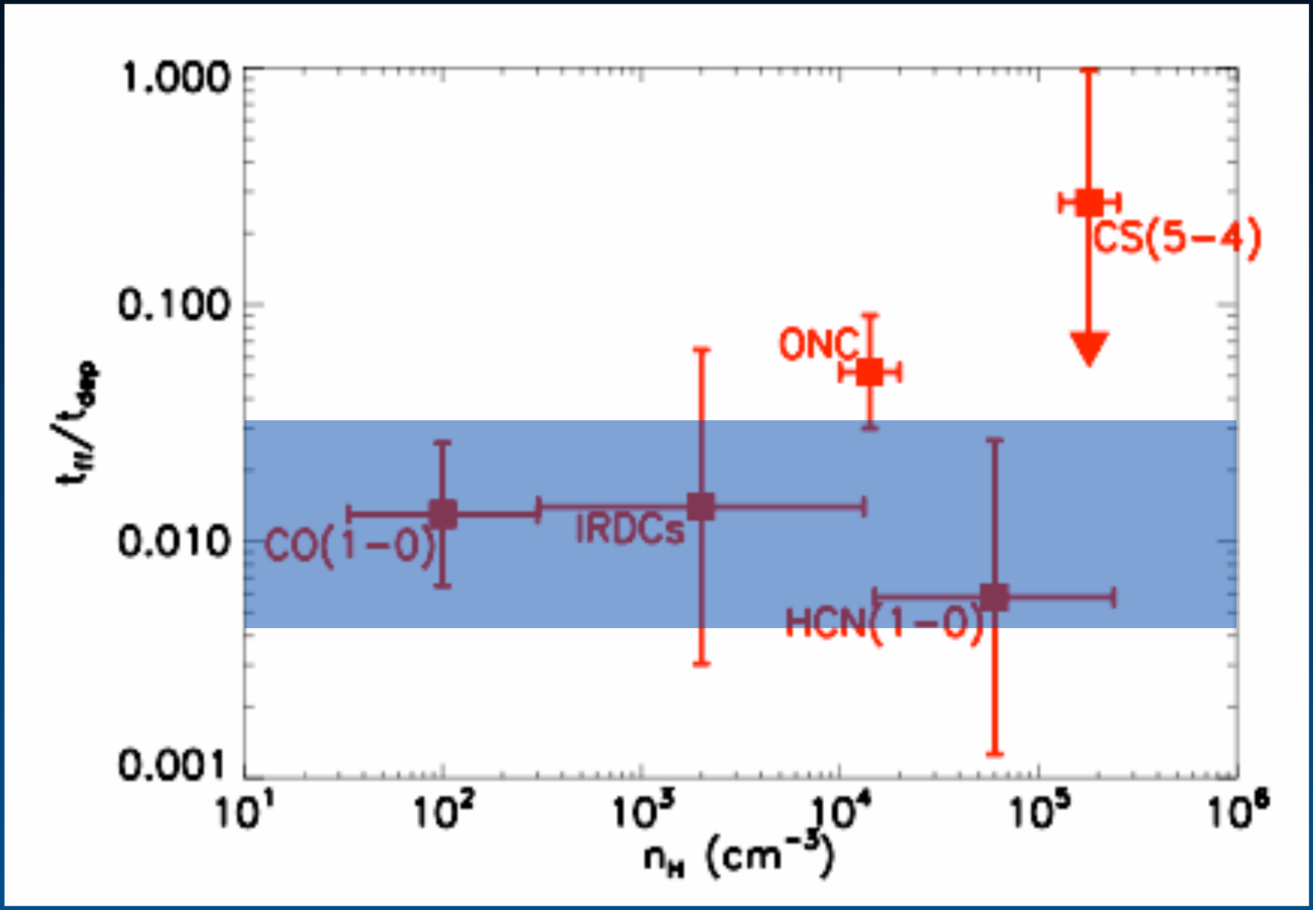
…even in dense gas



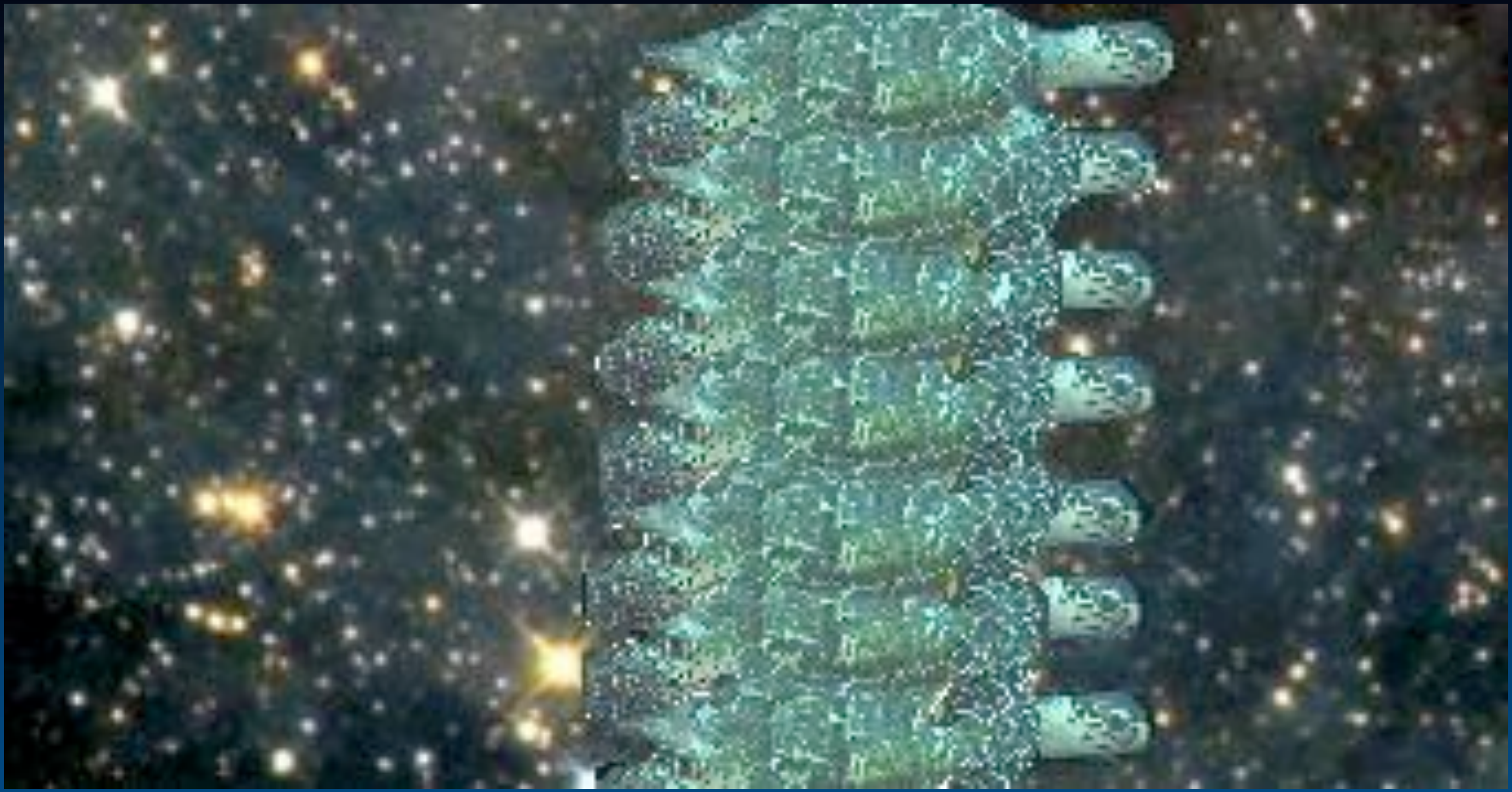


Depletion time as a function of Σ_{H2} for 2 local galaxies (left, Wong & Blitz 2002) and as a function of L_{HCN} for a sample of local and z ~ 2 galaxies (right, Gao & Solomon 2004, Gao et al. 2007)

There is a Universal SFR



Clouds convert ε_{ff} ~1% of their mass to stars per t_{ff}, regardless of density or environment (Tan, Krumholz, & McKee 2006; Krumholz & Tan 2007; Evans et al. 2009)



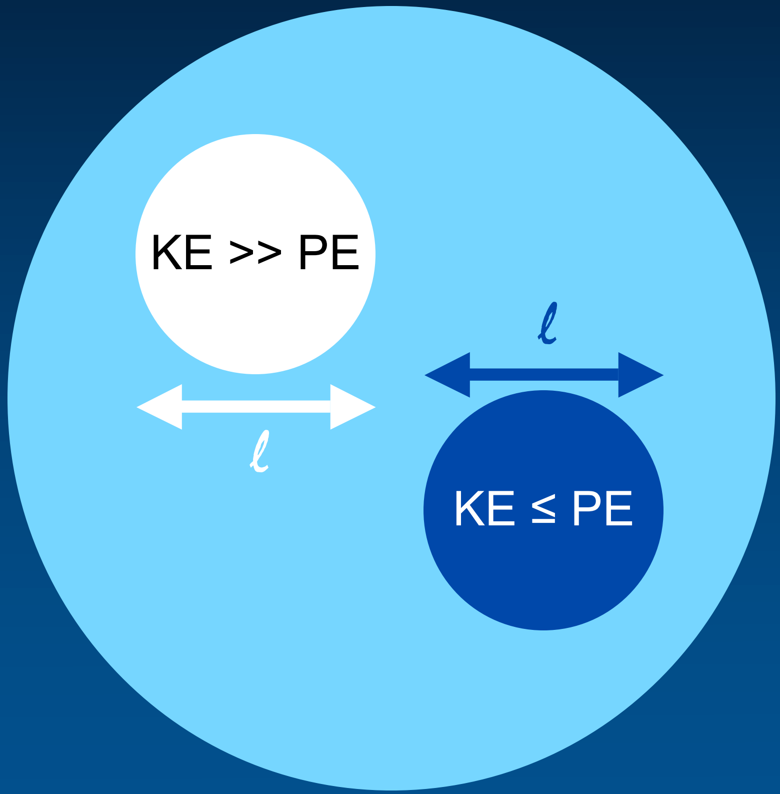
In other words:

it's turtles all the way down…

Where Does ε_{ff} Come From?

(Krumholz & McKee 2005)

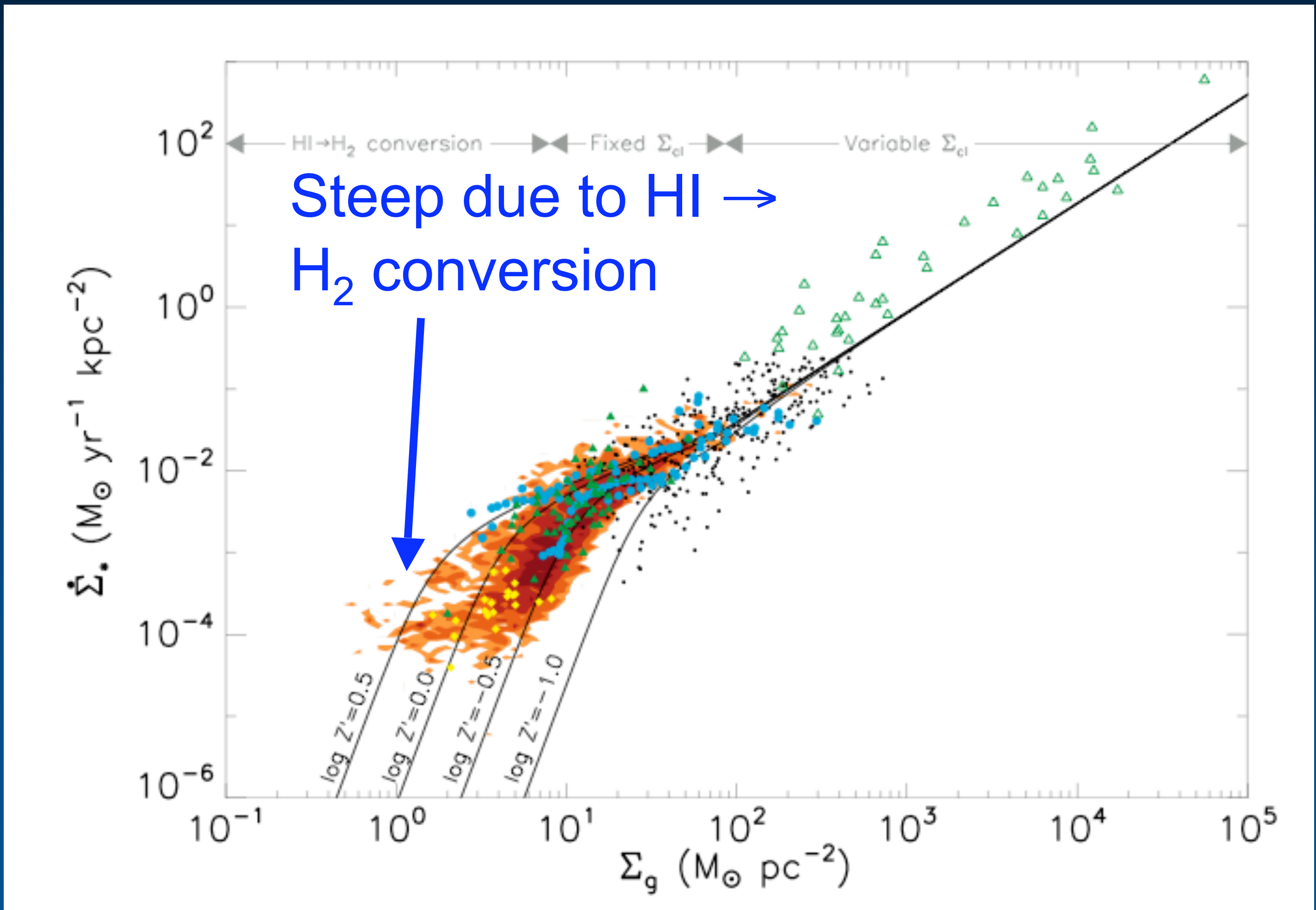
- On large scales, GMCs have $\alpha \approx 1$ (i.e. PE $\approx$ KE)
- Linewidth-size relation: $\sigma_v \approx c_s (\ell / \lambda_s)^{1/2}$
- In average region, $M \propto \ell^3$
  $\Rightarrow$ KE $\propto \ell^4$, PE $\propto \ell^5$
  $\Rightarrow$ KE >> PE
- SF occurs in overdense regions where PE ≥ KE
- Density PDF is lognormal
- Fraction of density with PE ≥ KE is ~1%



ε_{ff} ~ 1% for any turbulent, virialized object

Putting it Together: The Total Gas Star Formation Law

(Krumholz, McKee, & Tumlinson 2009)



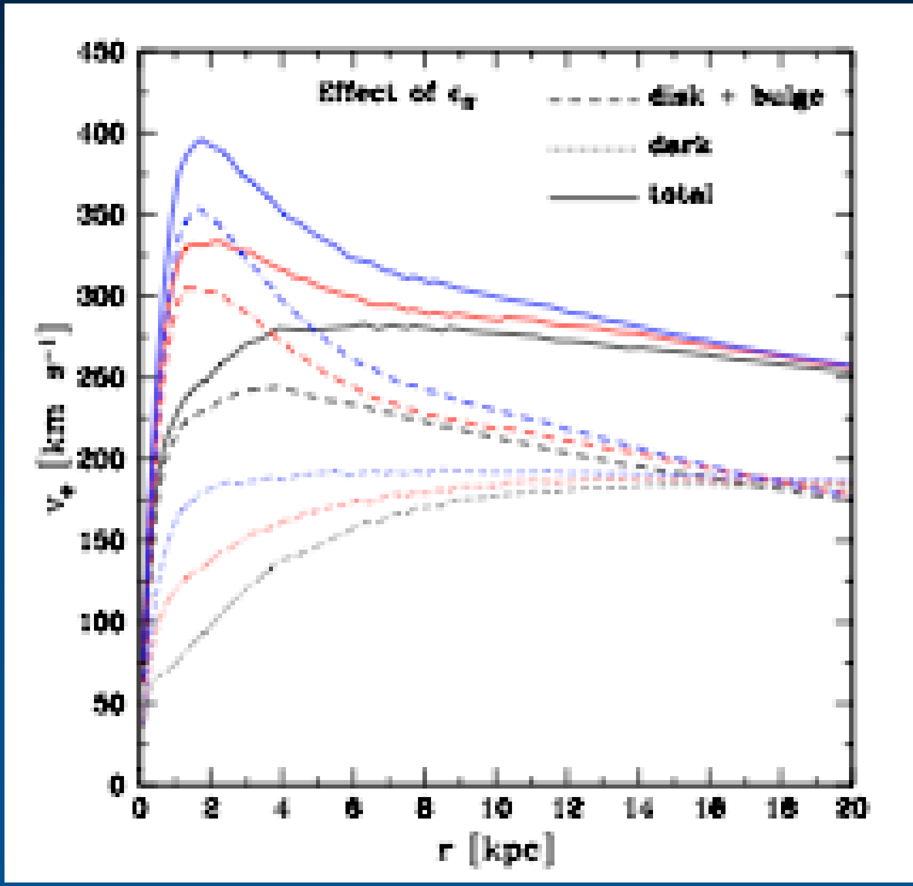
Lines: theory

Contours: THINGS, Bigiel et al. 2008

Symbols: literature data compiled by Bigiel et al. 2008

Suggested Implementation

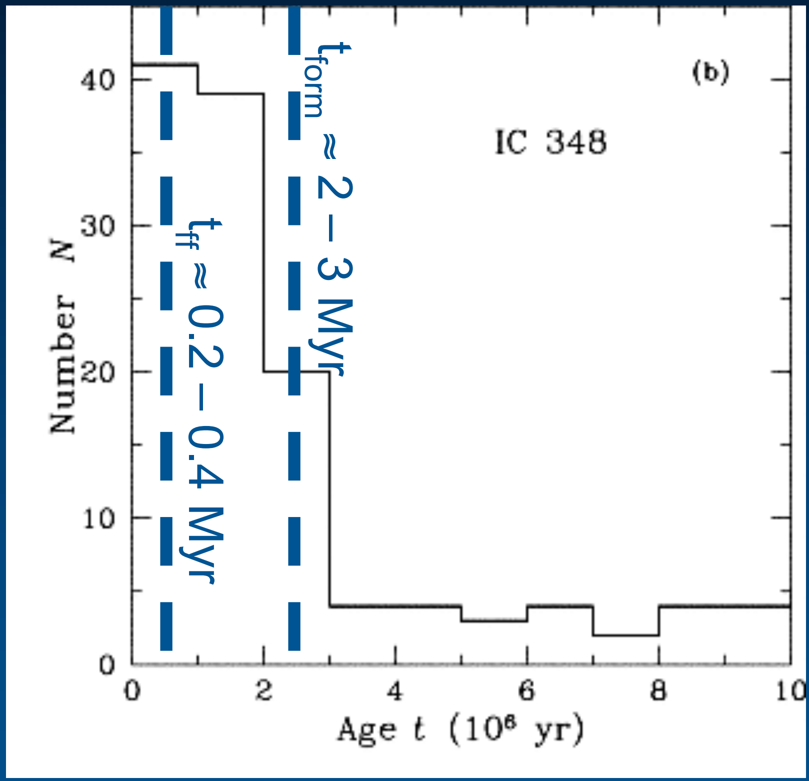
- Volumetric SFR $\dot{\rho}_* = f_{\mathrm{H_2}} \epsilon_{\mathrm{ff}} \frac{\rho}{t_{\mathrm{ff}}}, \; \epsilon_{\mathrm{ff}} = 0.01$



v_{circ} in a z = 0 galaxy simulated with ε_{ff} = 0.01 (black), 0.02 (red), 0.05 (blue); only ε_{ff} = 0.01 produces a flat rotation curve (Agertz, Teyssier, & Moore 2010)

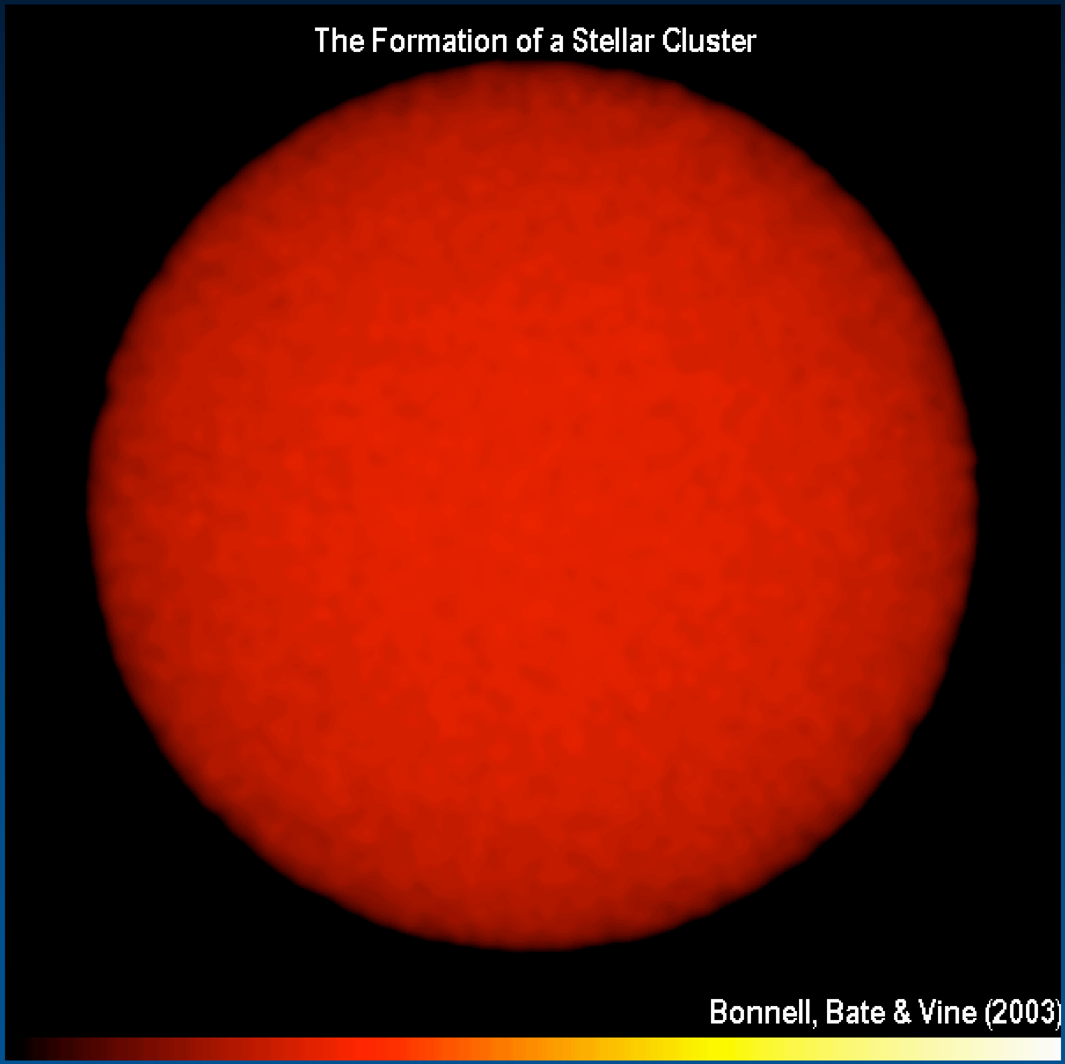
Feedback: More than Just Supernovae

- In star clusters forming today, the action is over before the first SN goes off
- Even in GMCs (~30 Myr lifetime), SN occur only after stars leave the cloud

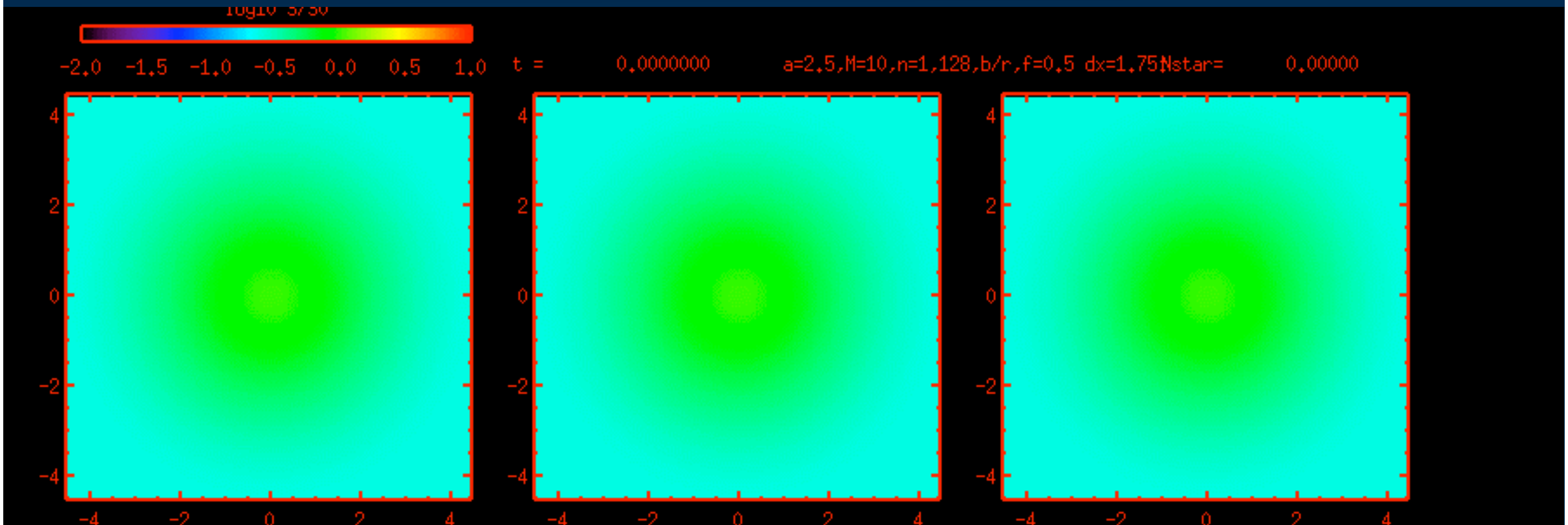


Numbers of stars in IC 348 vs. age inferred by pre-main sequence models, Palla & Stahler 2001

Cluster Formation without Feedback

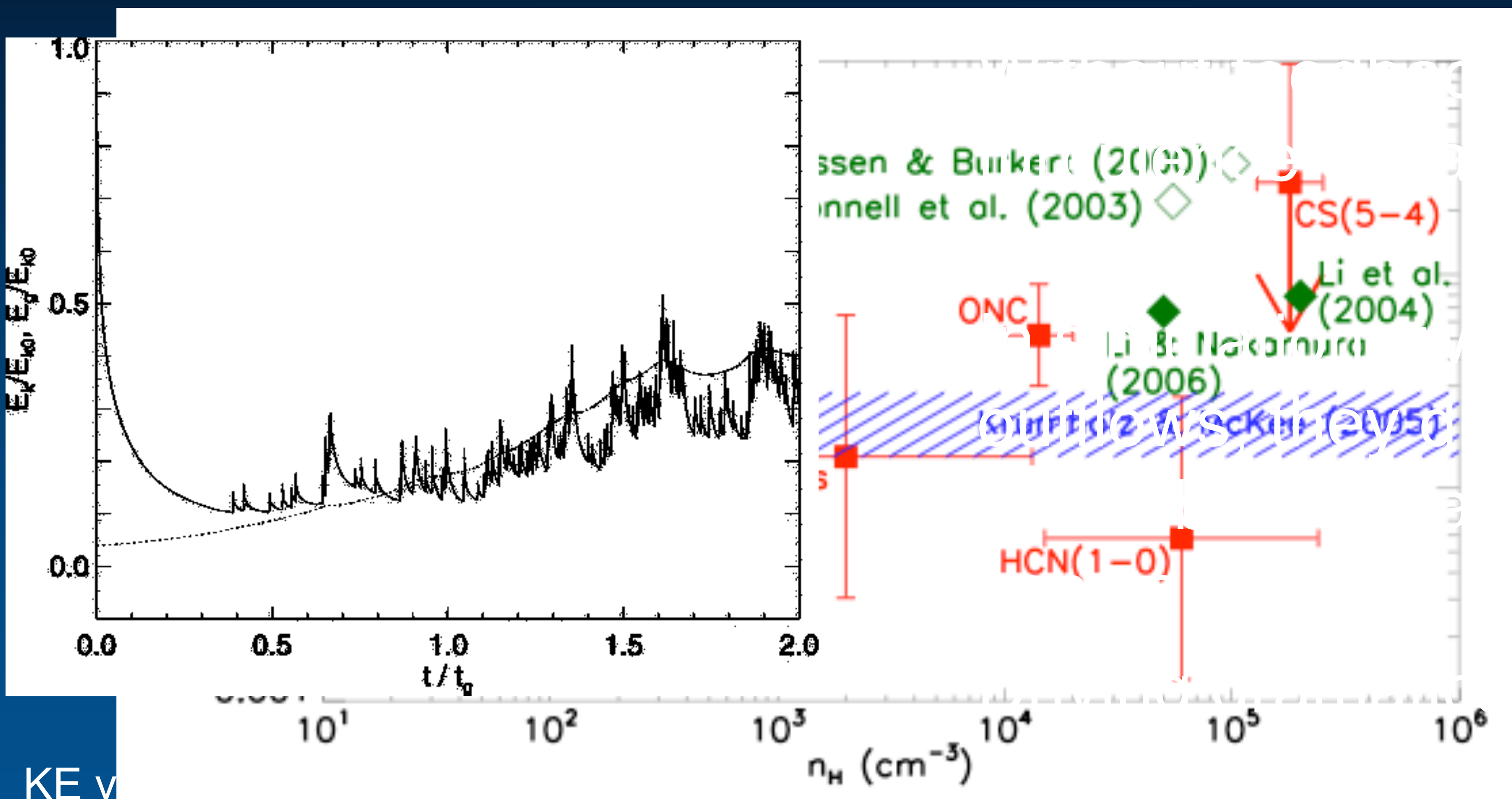


With Protostellar Outflows



Li & Nakamura (2006)

Outflows Drive Turbulence



KE v
protostellar outflows (Li & Nakamura 2006; also Wang+ 2010)

Comparison of observed and simulated SFRs (Krumholz & Tan 2007)

Ionization Feedback

30 Doradus HII region, MCELS

- On GMC scales, outflows probably cannot drive turbulence (Matzner 2002, 2007)
- Observed GMC lifetime is ~ 30 Myr (Blitz et al. 2007), t_{ff} ~ 4 Myr ⇒ turbulence must be driven
- HII regions are the most likely candidate

Simulation of HII Region Feedback

(Krumholz, Stone, & Gardiner 2007; see also Grittschneder et al. 2009)

A Semi-Analytic GMC Model

Follow evolution of:
M_{gas}, M_*,
R, dR/dt, σ

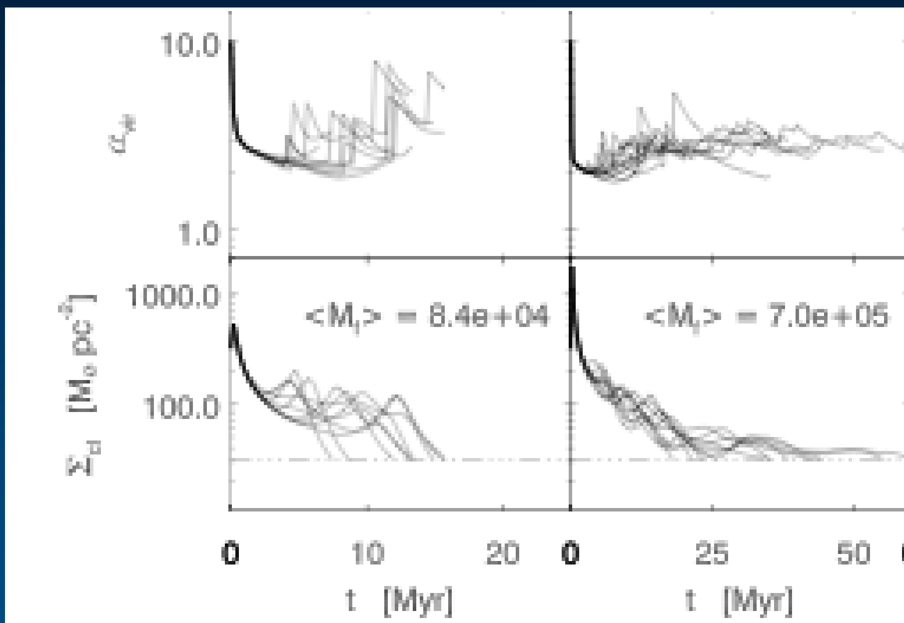
- Model GMC mass, energy, momentum budgets, with feedback and mass loss
- Evolution controlled by energy and virial equations, with energy injection by HII regions and loss by decay of turbulence:

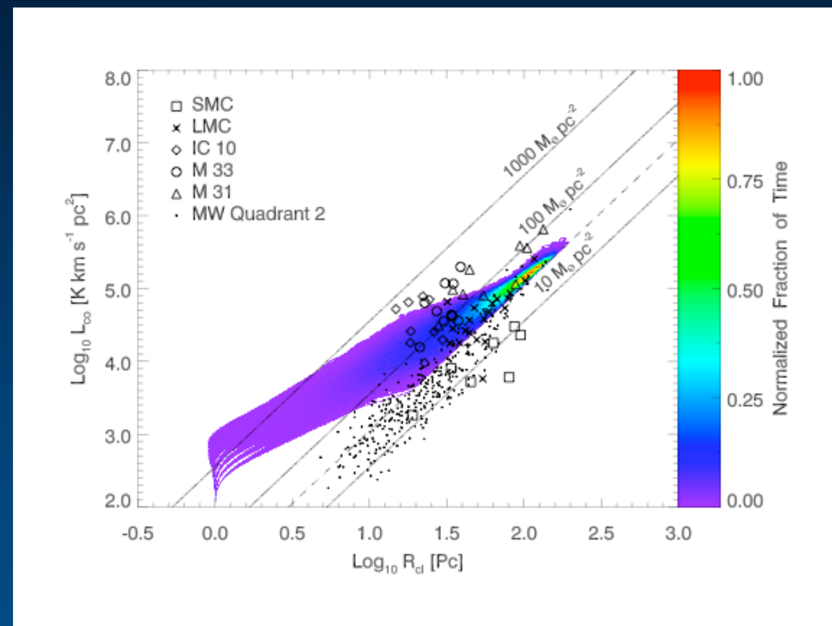
$$\frac{\ddot{I}}{2} = 2(\mathcal{T} - \mathcal{T}_0) + \mathcal{W} + \mathcal{B} - \left(\frac{1}{2}\right) \frac{d}{dt} \int (\rho \mathbf{v} r^2) \cdot d\mathbf{S}$$

$$\dot{E} + \int \rho \left(\frac{v^2}{2} + e + \phi + \frac{P}{\rho} \right) \mathbf{v} \cdot d\mathbf{S} = \Gamma - \Lambda$$

Results of Ionization Feedback

(Goldbaum+ 2010, in preparation)



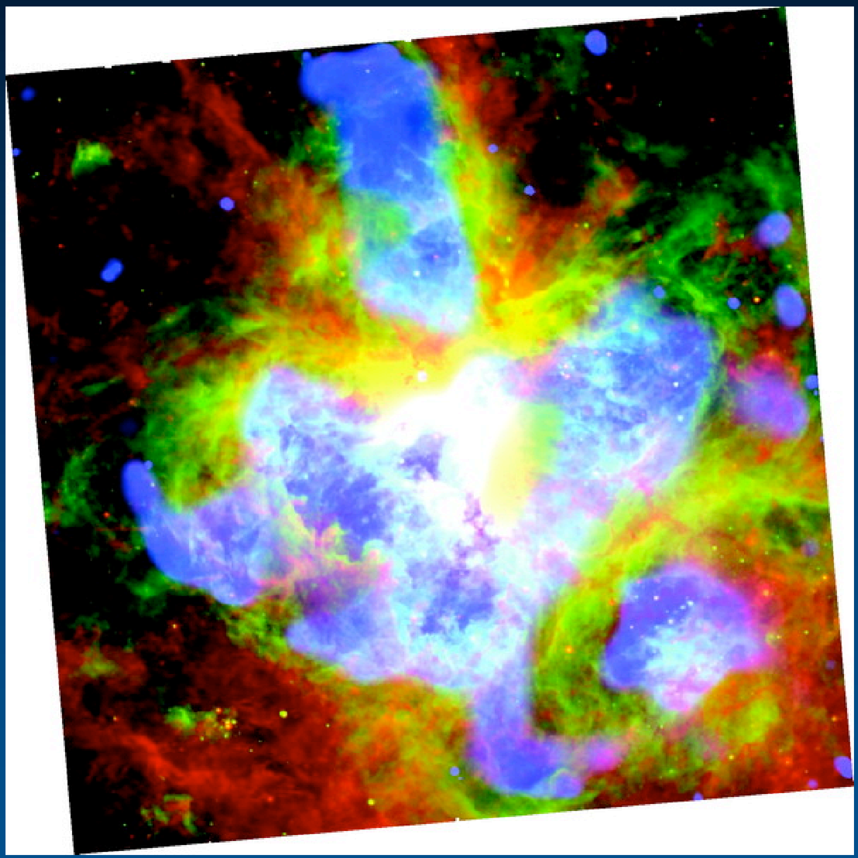


Left: GMC virial ratio and surface density versus time; at time of dissociation, total SFE is ~few percent

Right: GMC mass-radius distribution

Feedback at High Mass, Σ

(Krumholz & Matzner 2009; Murray, Quataert, & Thompson 2009)



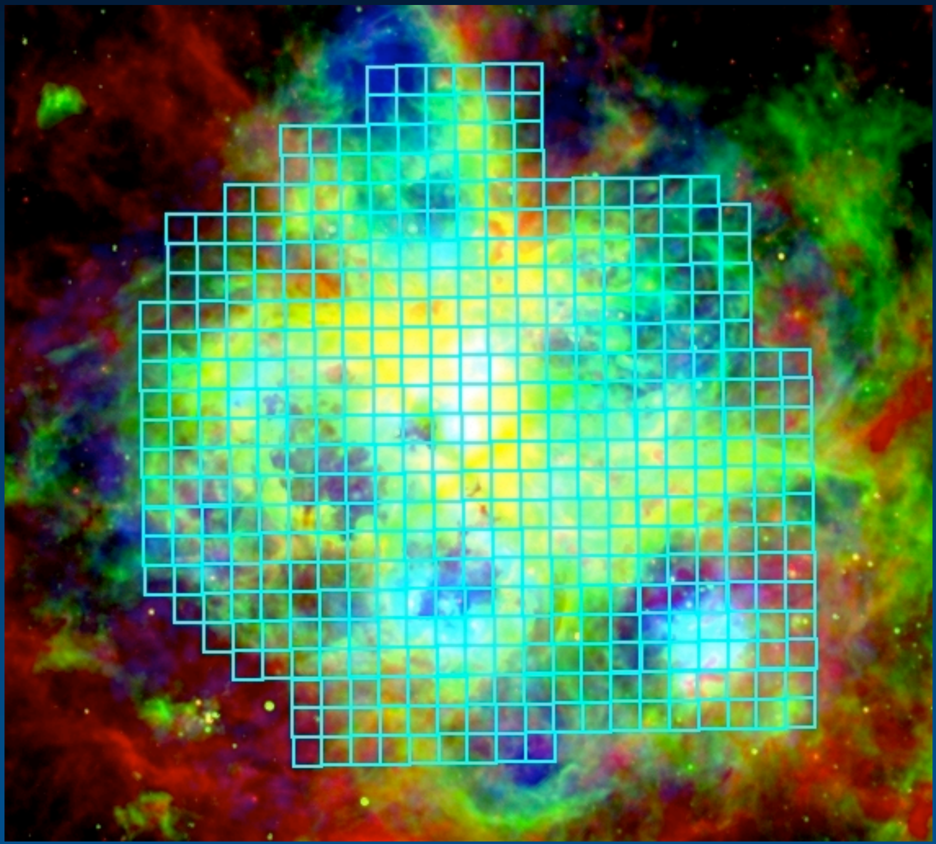
30 Doradus HII region in IR (red), Hα (green), x-ray (blue) (Townsley+ 2006)

- For massive proto-clusters, ionized gas pressure is ineffective
  - Ex.: R136, $M = 5 \times 10^4$ $M_\odot$, R = 1 pc, $v_{esc} = 20$ km s^{-1} $\approx 2\, c_{II}$
- SNe are too late
  - For R136, $t_{cr} \sim 50$ kyr
- Only possibilities: hot gas from winds, radiation pressure

Observational Test: 30 Doradus

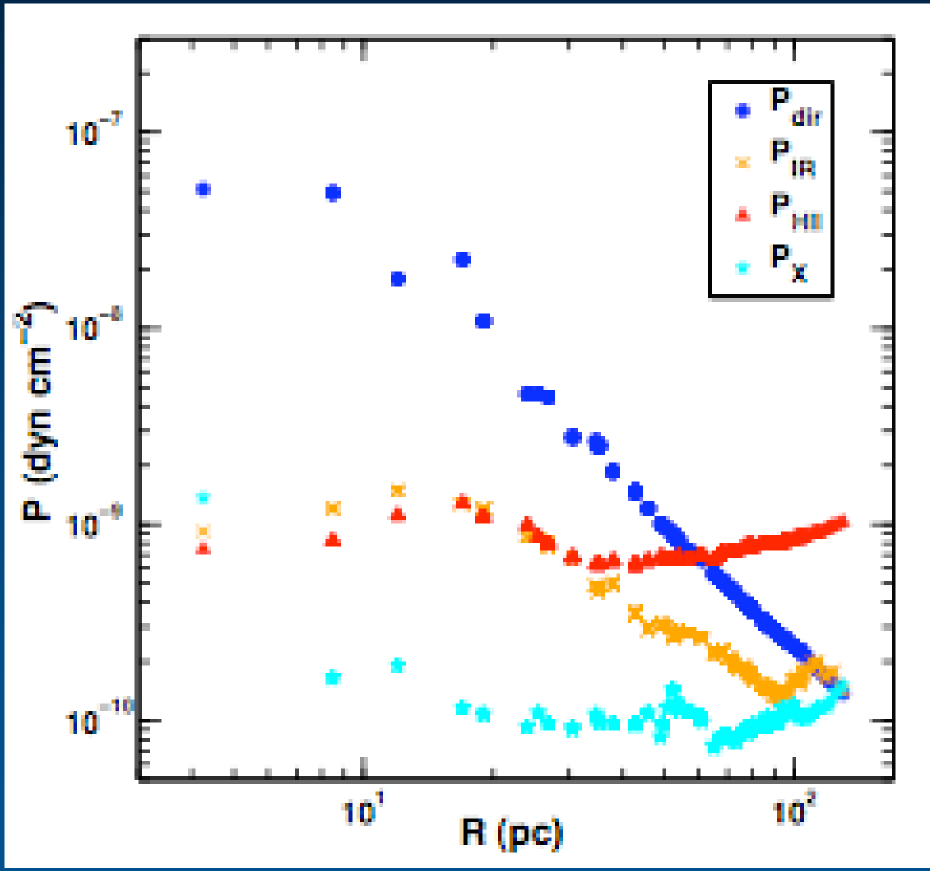
(Lopez+ 2010, in preparation)

- Measure sources of pressure in 30 Dor
  - Hot shocked gas (wind, SNe): x-rays / Chandra
  - Warm ionized gas: radio continuum
  - Direct radiation pressure: optical
  - Dust-processed rad. pressure: IR / Spitzer
- Compute pressure pixel-by-pixel



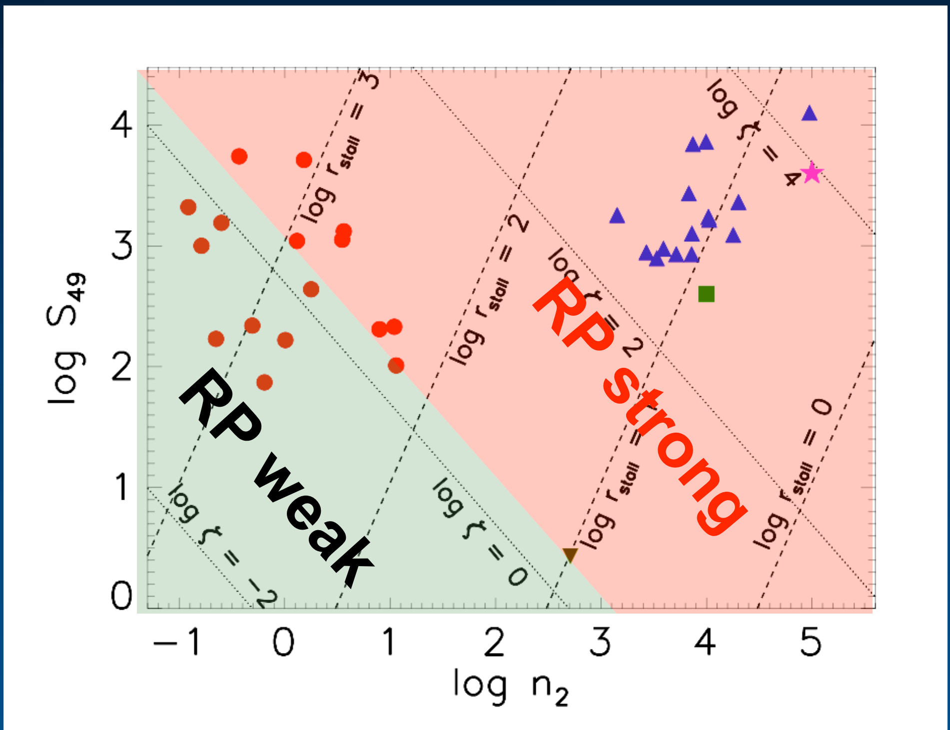
30 Doradus overlayed with pixels for pressure computation (blue = x-ray, green = Hα, red = 8 μm)

Direct Radiation and Warm Gas Dominate the Pressure



When is Radiation Pressure Important in HII Regions?

(Krumholz & Matzner 2009)



Importance of RP in clusters in M82 (blue), Antennae (red), Orion (brown), Arches (green)

- RP force >> gas pressure force when

$$\zeta = 6.2 \times 10^{-2} n_2^{2/3} S_{49}^{2/3} \gg 1$$

- RP-driven expansion stalls at radius

$$r_{\rm st} = 8.9 n_2^{-1/2} S_{49}^{1/4} \;\mathrm{pc}$$

- Ex. R136: $n_2 \sim 10^3$, $S_{49} \sim 10^2 \Rightarrow \zeta \sim 100$, $r_{st} \sim 1$ pc

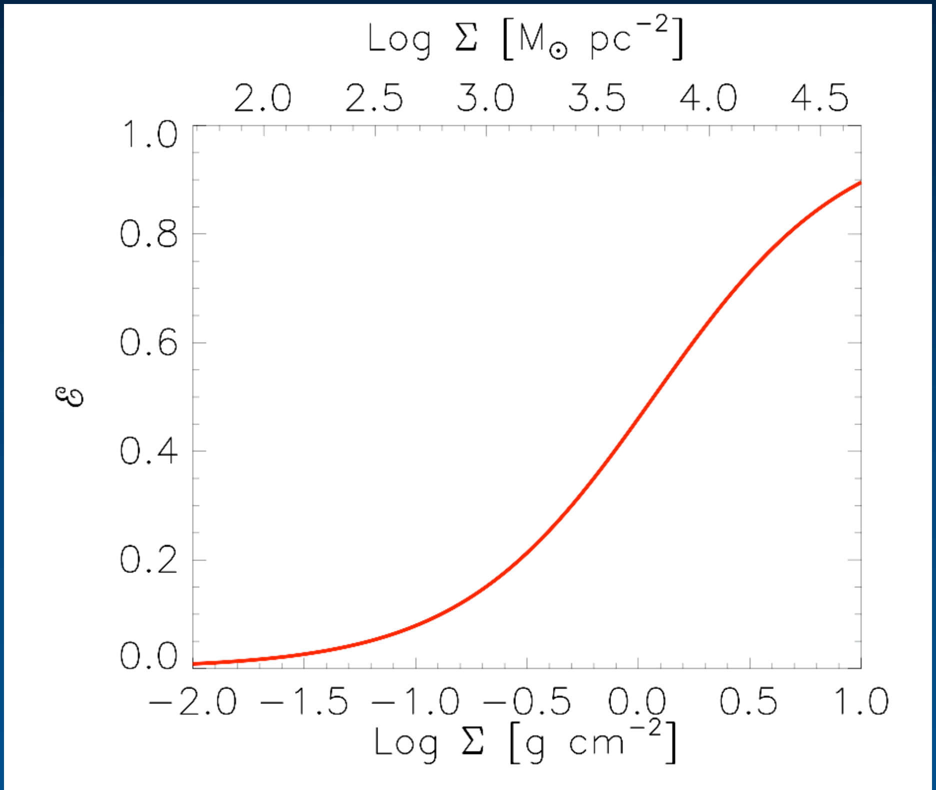
Star Formation Efficiency from Radiation Pressure

(Fall, Krumholz, & Matzner 2010)

- Rough SFE estimate: as SF proceeds and SFE rises, n_2 drops, r_{st} rises
- When $r_{st} > R_{cl}$, mass is ejected
- Result:

$$\mathcal{E} \simeq \frac{\Sigma}{\Sigma + 1.2 \text{ g cm}^{-2}}$$

- NB: depends on M only through Σ

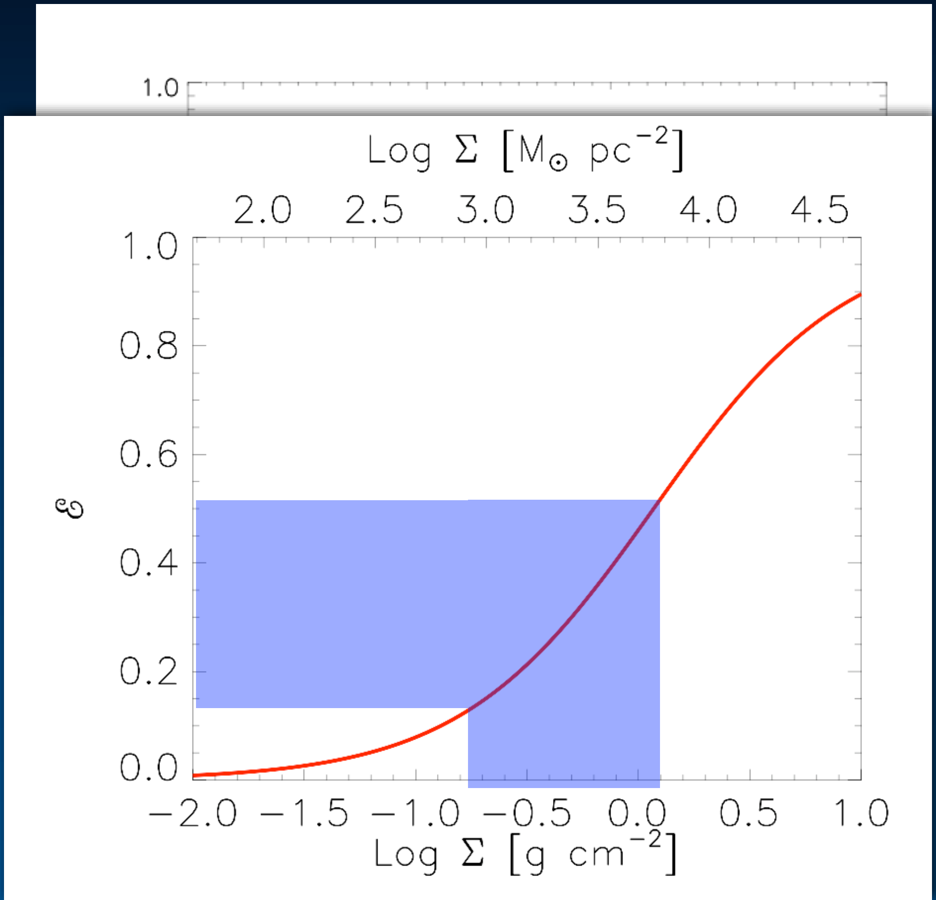


SFE vs. Σ, computed using RP feedback

Implications for Cluster MF

(Fall, Krumholz, & Matzner 2010)

- Protoclusters have Σ ~ 0.2 - 1 g cm^{-2} independent of M ⇒ SFE independent of M
- For observed Σ, SFE ~ 0.2 - 0.4 ⇒ most but not all clusters dissolve at all M
- Cluster MF ~ same as cloud MF, in agreement with observations



Cluster-forming clumps: CS emission (Shirley+ 2003, black), dust (Faundez+ 2004, blue), C^{17}O (Fontani+ 2005, red)

Feedback Beyond SNe

- At scales below ~100 pc, non-SN feedback is required to produce correct SF rate, efficiency, cluster distribution
- Numerical implementation depends on scale and type of feedback
- Subgrid models probably preferable to direct simulation unless resolution is very high (~1 pc or better)